FINDING THE FIRST METALS

RYAN JAMES COOKE

INSTITUTE OF ASTRONOMY &
TRINITY COLLEGE

This dissertation is submitted for the degree of Doctor of Philosophy

May 2011
DECLARATION

I hereby declare that my thesis entitled *Finding the First Metals* is not substantially the same as any that I have submitted for a degree or diploma or other qualification at any other University. I further state that no part of my thesis has already been, or is being concurrently submitted, for any such degree, diploma or other qualification. The containing work was completed during my candidature for the degree of Doctor of Philosophy at the University of Cambridge.

Furthermore, the work presented in this dissertation is my own and contains nothing which is the outcome of work done in collaboration with others, except as specifically indicated clearly in the text and acknowledgements. Whilst all writing was generated by myself, I have asked others (credited in the acknowledgements and in Section 2.5) for their aid in editing and revision. Those parts of this thesis which have been published or accepted for publication are as follows.

- Chapter 3 was published as:
  

- Chapter 5 was published as:
  

- Chapter 7 was published as:
  

Various figures throughout the text are reproduced from the work of other authors, for illustration or discussion. Such figures are always credited in the associated caption.

This thesis contains fewer than 60,000 words.

R. J. Cooke
Cambridge, 9 May 2011
SUMMARY

Our knowledge of the birth, life, and death of the objects that commenced the epoch of reionization is far from complete. We do however know that these objects, the first stars, played a crucial role in creating and distributing the first metals in our Universe, hence setting the initial conditions for cosmic chemical evolution. Until recently, our understanding of the first generation of stars largely came from observations of the most metal-poor stars in the halo of the Milky Way, which are believed to have condensed out of pristine gas that was solely enriched by the first generation. In this thesis I explore a new avenue to study the nature of the first stars.

In Chapter 1, I review our present understanding of the formation, evolution, and death of the first stars, from both a theoretical and observational perspective. I also include a summary of the model yield calculations from metal-free stars and comment on how these first metals assist in the transition from the first generation of massive metal-free stars to the low-mass dominated population of extremely metal-poor stars that we observe in the Universe today.

The underlying goal of this thesis is to uncover the clouds of gas that gave birth to the second generation of stars. I outline the expected properties of such clouds in Chapter 2, and suggest that these properties are satisfied by the most metal-poor damped Ly$\alpha$ systems. I conclude this Chapter by presenting the current evidence in support of this idea, and summarising the methods that I will use in this thesis to explore these near-pristine clouds of gas.

Chapter 3 reports a previously unidentified damped Ly$\alpha$ system along one sightline towards the gravitationally lensed quasar pair UM 673. This example, together with other examples in the literature, provides an estimate for the typical sizes of damped Ly$\alpha$ systems. Along the neighbouring sightline, I detect a weak Ly$\alpha$ emission line that I attribute to nearby star formation.

In Chapter 4, I present the results from my study of the chemical abundances in the most metal-poor damped Ly$\alpha$ systems. I confirm the recent suggestion that stars and damped Ly$\alpha$ systems exhibit an elevated [C/O] abundance in this regime, and provide new light on the much-debated nature of the [O/Fe] ratio at low metallicity. Finally, I examine the possibility that these systems contain the nucleosynthetic products of the first stars, but a definite conclusion is premature at this stage. Additional support in favour of this idea is presented in Chapter 5, where I report the discovery of an extremely metal-poor damped Ly$\alpha$ system that exhibits the nucleosynthetic imprint of the first stars. One of the key signatures identified is a [C/Fe] ratio of $+1.53$, unlike any other damped Ly$\alpha$ system known to date, yet in good agreement with some of the carbon-enhanced metal-poor stars in the halo of our Galaxy.

Finally, in Chapter 6 I outline the important conclusions that have resulted from this work, and suggest several element abundance ratios that could help to identify the systems that are the most likely to exhibit the chemical signature of the first stars. To conclude, I propose a means to estimate the typical explosion energy of the stars that polluted these systems with their metals.

To supplement the work in this thesis, I include additional, unrelated work in Chapter 7 that searches for a systematic dipole in the distance estimates from type Ia supernovae.
To those that made this possible.
ACKNOWLEDGEMENTS

Somehow, a thesis follows on from this page. This would not have been possible without the generous support of my funding bodies, including the Cambridge Overseas Trust and the Cambridge Commonwealth/Australia Trust. I must personally thank Charles Allen, for the award of an Allen Cambridge Australia Trust Scholarship. Your generosity far outweighs anything I could repay. Thank you for giving me this opportunity. In addition, I would also like to thank Trinity College for the additional monetary support, accommodation, and the (sometimes) delicious meals that they provided.

The majority of the work presented in this thesis was completed with the continuing guidance and support (not to mention tolerance!!!) from my primary supervisor, [Max Pettini]. The square brackets around your name indicate, as in the “Select a Supervisor” manual given to me on arrival at the IoA, that you were not likely to take on a new Ph.D. student in my year due to a sabbatical leave. I ‘apologise’ for finding your work so interesting (and for asking you to supervise me), but I must say, you have gone beyond my expectations for a supervisor and I could not possibly have made a better choice. Thank you for all of your assistance that has resulted in the completion of this thesis, not to mention the games of cricket and the ‘barbies’ at your place. I am also indebted to Donald Lynden-Bell, my secondary supervisor, who advised me at the start of my Ph.D., and showed continued interest in my work. I also thank you for the many fruitful discussions we had about things in the Universe... on various physical scales!

I also wish to thank my collaborators, Chuck Steidel, Gwen Rudie, Poul Nissen, Lindsay King, Regina Jorgenson, and Olivera Rakic, who assisted me with the work presented herein. I also benefited from the fruitful discussions that I had with many others, including George Becker, Sebastiano Cantalupo, Bob Carswell, Martin Haehnelt, Chiaki Kobayashi, Geraint Lewis, Eric Monier, Sam Rix, Nozomu Tominaga, and Chris Tout. A special thank you goes to Paul Hewett who provided additional guidance along the way (on jobs, science, 1st, 2nd, and 3rd year reports... the list goes on). Furthermore, throughout I have used the freely available software developed by others to perform difficult tasks in a fraction of the time it would take me to perform them, including the software developed by Tom Barlow, George Becker, Bob Carswell, Gary Ferland, Alexander Heger, Robert Lupton, Patricia Monger, and Michael Murphy. Facilities, such as European Southern Observatory and the Keck Observatory, provided most of the data presented in this thesis.

In addition, I thank Bob Carswell and Francesca Primas, my thesis examiners, for (passing me and) providing useful comments that I have incorporated into my thesis.

I thank my past and present office mates, Kostas, Damien, Mark, James, Chia-Ying, and Steve, for providing an enjoyable working environment and for coping with the incessant clicking of my mouse. To my year group (who, for some reason, I still refer to as “the first years”), Warrick, Amy, Jon, Alex, Becky, Stephanie, Yin-Zhe, James M., James O., Adrian, and Dom, thank you for providing barrel’s of laughter and, at times, (semi-)serious discussions.
At first, I hesitated to accept my offer into Cambridge, and if it wasn’t for the gentle encouragement by Zdenka Kuncic and Joss Bland-Hawthorn, I may not be writing this acknowledgement today. I must also thank the lovely administration team at the IoA, whom I asked many questions to and always received an answer. A special thanks goes to Siân Owen for simply being wonderful and supportive no matter what the circumstance.

Another special thank you goes to my parents and family, for their endless and ongoing encouragement. Thank you for always being there for me on email or on Skype when I needed advice or simply wanted to chat. To Nathaniel, you’ll understand where I am when you’re a little bit older.

And, finally, to Alis – my heart throbs even now when I write your name – you are the love of my life, my wife-to-be, and more supportive than I could ever imagine. Thank you so much, for just everything.

Ryan J. Cooke
Cambridge, 9 May 2011

This thesis was typeset in \LaTeX using GNOME Text Editor.
## Contents

Declaration iii  
Summary v  
Acknowledgements ix  
Contents xi  
List of Figures xiv  
List of Tables xvii

### 1 The First Stars
1.1 The birth, life, and death of the first stars 1  
1.2 The nucleosynthesis of the first metals 8  
1.3 The transition to Population II star formation 13  
1.4 Finding the first metals: Chemical signatures from the Galactic halo 15

### 2 The Missing Link
2.1 Damped Lyman-α systems 19  
2.2 Host Galaxy Identification 21  
2.3 DLAs as Chemical Laboratories 24  
2.3.1 Absorption line formation 24  
2.3.2 Observational techniques and analysis 27  
2.3.3 Studying the chemical properties of DLAs 31  
2.4 Finding the first metals: Chemical signatures from DLAs 34  
2.5 Organisational Notes 39

### 3 The sizes and star formation rates of DLAs
3.1 Introduction 41  
3.2 Observations and Data Reduction 43  
3.3 H I Absorption towards UM 673A,B 44  
3.3.1 The DLA towards UM 673A 44  
3.3.2 The Lyman limit system in UM 673B 45  
3.4 Constraining the sizes of DLAs 47  
3.5 Lyα emission towards UM 673B 52
3.5.1 Origin of the Ly\(\alpha\) emission .......................... 53
3.5.2 Lensed Ly\(\alpha\) emission? .............................. 54
3.6 Chemical Composition of the DLA in front of UM 673A .......................... 55
   3.6.1 Profile Fitting ........................................ 55
   3.6.2 Element Abundances .................................. 61
3.7 Summary and Conclusions ..................................... 67

4 A survey for the most metal-poor DLAs .......................... 69
   4.1 Introduction .............................................. 70
   4.2 Observations and Data Reduction ............................. 72
      4.2.1 Target Selection .................................... 72
      4.2.2 Echelle spectroscopic follow-up ....................... 73
      4.2.3 Data Reduction ..................................... 76
   4.3 Profile Fitting ........................................... 77
   4.4 Individual Objects ....................................... 79
      4.4.1 J0311$-$1722: DLA at $z_{\text{abs}} = 3.73400$ .......................... 79
      4.4.2 J0831+3358: DLA at $z_{\text{abs}} = 2.30364$ .......................... 81
      4.4.3 J1001+0343: DLA at $z_{\text{abs}} = 3.07841$ .......................... 81
      4.4.4 J1037+0139: DLA at $z_{\text{abs}} = 2.70487$ .......................... 84
      4.4.5 J1340+1106: DLA at $z_{\text{abs}} = 2.50792$ .......................... 86
      4.4.6 J1340+1106: DLA at $z_{\text{abs}} = 2.79583$ .......................... 88
      4.4.7 J1419+0829: DLA at $z_{\text{abs}} = 3.04973$ .......................... 91
      4.4.8 The final VMP DLA sample ................................ 91
   4.5 Abundance Analysis ....................................... 93
      4.5.1 Ionisation Corrections ................................ 93
      4.5.2 Dust Depletion ...................................... 96
   4.6 Comparing VMP DLAs and stars ............................... 96
      4.6.1 Revisiting C/O at low metallicity ....................... 97
      4.6.2 The O/Fe debate in the metal-poor regime ............... 100
   4.7 Discussion ............................................... 103
      4.7.1 The typical VMP DLA ................................ 103
      4.7.2 Clues to early episodes of nucleosynthesis ............... 105
      4.7.3 Comparison with data of medium spectral resolution ....... 108
   4.8 Summary and Conclusions .................................. 110

5 The nucleosynthesis from the first stars ......................... 113
   5.1 Introduction .............................................. 114
   5.2 Observations and Data Reduction ............................. 117
   5.3 Profile Fitting and Abundance Analysis ......................... 118
5.3.1 Column Densities ................................................. 120
5.3.2 Ionisation Corrections ........................................... 121
5.4 A masquerading carbon enhancement? ......................... 123
  5.4.1 Incorrect background subtraction? ......................... 123
  5.4.2 Profile fitting 1: gas kinematics ......................... 123
  5.4.3 Profile fitting 2: instrumental resolution ............... 124
  5.4.4 Profile fitting 3: thermal broadening of the line profiles 124
  5.4.5 Monte Carlo simulations .................................. 126
5.5 A Carbon-Enhanced metal-poor DLA ............................ 128
  5.5.1 Comparison with Stellar Abundances .................... 130
  5.5.2 Comparison with Model Yields for Metal-free Stars ........ 132
  5.5.3 How Many Supernovae? .................................. 134
5.6 Summary and Conclusions ....................................... 135

6 Finding the First Metals ............................................ 137
  6.1 What fraction of DLAs are carbon-enhanced? ................ 138
  6.2 Uncovering the nature of the first stars .................... 142

7 Does the Universe expand (an)isotropically? ........................ 147
  7.1 Introduction ................................................... 148
  7.2 The Type Ia Supernova Sample ............................... 149
  7.3 Analysis ..................................................... 150
    7.3.1 The Hubble Deviation ................................. 150
    7.3.2 Maximum Likelihood Strategy ......................... 151
  7.4 Results and Discussion ..................................... 154
    7.4.1 Significance of the Dipolar Models .................. 154
    7.4.2 The Weighted Dipolar Model ......................... 155
    7.4.3 The Higher Quality List .............................. 156
  7.5 Conclusions ............................................... 157
  7.6 Future Work ............................................... 158

APPENDICES

A The adopted Solar abundance scale ................................ 163

B Column densities for VMP DLAs .................................. 165

C [O/Fe] measurements for VMP DLAs & stars ...................... 167

References .......................................................... 169
List of Figures

1.1 The properties of a collapsing cloud of primordial gas. 3
1.2 The formation of the first stars in a cosmological simulation. 5
1.3 Pop III small multiples in a primordial protostellar disc. 6
1.4 The initial-final mass function for non-rotating metal-free stars. 7
1.5 The mixing-fallback model of metal-free nucleosynthesis calculations. 9
1.6 The effect of Rayleigh-Taylor induced mixing in metal-free stars. 11
1.7 The metal yields from core-collapse and pair-instability supernovae of metal-free stars. 12
1.8 The transition discriminant for low mass star formation 14
1.9 Fitting the lowest metallicity halo stars to Pop III yields. 16
1.10 Comparing the C/O ratio in the most metal-poor stars with Galactic chemical evolution models, considering only Population II nucleosynthesis. 17

2.1 A demonstration of Quasar absorption line spectroscopy. 20
2.2 An example quasar spectrum that contains a DLA. 20
2.3 An image of the suspected host galaxy of a DLA. 22
2.4 DLAs in a cosmological simulation of galaxy formation. 23
2.5 Example Voigt profiles from the linear, flat, and damped regimes. 25
2.6 The curve of growth. 26
2.7 A schematic of the HIRES instrument. 28
2.8 The echelle data format for the HIRES detector. 29
2.9 The metallicity and dust-depletion distributions for DLAs. 33
2.10 The observed and simulated metallicity distribution of DLAs. 34
2.11 The mass-metallicity relation in DLAs. 35
2.12 Comparing a DLA’s abundance pattern to model yield calculations of the first stars. 36
2.13 The C/O ratio in the most metal-poor stars and DLAs (high resolution). 37
2.14 The C/O ratio in the most metal-poor stars and DLAs (low resolution). 38

3.1 An image of the lensed quasar UM 673 42
3.2 The Lyα line for the DLA towards UM 673A 45
<table>
<thead>
<tr>
<th>Figure</th>
<th>Description</th>
</tr>
</thead>
<tbody>
<tr>
<td>3.3</td>
<td>A selection of absorption lines in UM 673B near the redshift of the DLA in front of UM 673A</td>
</tr>
<tr>
<td>3.4</td>
<td>The geometry of the gravitational lens</td>
</tr>
<tr>
<td>3.5</td>
<td>The distributions for the derived scale lengths of DLAs</td>
</tr>
<tr>
<td>3.6</td>
<td>SDSS spectrum of UM 673A,B</td>
</tr>
<tr>
<td>3.7</td>
<td>A selection of metal absorption lines associated with the DLA in UM 673A</td>
</tr>
<tr>
<td>3.8</td>
<td>The weakest absorption features of the DLA are reproduced on an expanded scale</td>
</tr>
<tr>
<td>3.9</td>
<td>Transitions from highly ionised gas at redshifts close to that of the DLA in UM 673A</td>
</tr>
<tr>
<td>3.10</td>
<td>Ionisation corrections appropriate for the DLA towards UM 673A</td>
</tr>
<tr>
<td>3.11</td>
<td>Comparing the abundance pattern of the DLA towards UM 673A to a typical metal-poor DLA and star</td>
</tr>
<tr>
<td>3.12</td>
<td>The [Zn/Fe] ratio in DLAs, stars, and dwarf spheroidal galaxies</td>
</tr>
<tr>
<td>4.1</td>
<td>An example SDSS spectrum that contains a VMP DLA</td>
</tr>
<tr>
<td>4.2</td>
<td>Selected absorption lines for the DLA towards J0311−1722</td>
</tr>
<tr>
<td>4.3</td>
<td>Selected absorption lines for the DLA towards J0831+3358</td>
</tr>
<tr>
<td>4.4</td>
<td>Selected absorption lines for the DLA towards J1001+0343</td>
</tr>
<tr>
<td>4.5</td>
<td>Selected absorption lines for the DLA towards J1037+0139</td>
</tr>
<tr>
<td>4.6</td>
<td>Selected absorption lines for the DLA towards J1340+1106 at z_{abs} = 2.50792</td>
</tr>
<tr>
<td>4.7</td>
<td>Selected absorption lines for the DLA towards J1340+1106 at z_{abs} = 2.79583</td>
</tr>
<tr>
<td>4.8</td>
<td>Selected absorption lines for the DLA towards J1419+0829</td>
</tr>
<tr>
<td>4.9</td>
<td>Ionisation corrections applicable for very metal-poor DLAs</td>
</tr>
<tr>
<td>4.10</td>
<td>C and O abundances in stars and DLAs</td>
</tr>
<tr>
<td>4.11</td>
<td>The [O/Fe] ratio in very metal-poor stars and DLAs</td>
</tr>
<tr>
<td>4.12</td>
<td>Comparing the distribution of [O/Fe] values in stars and DLAs</td>
</tr>
<tr>
<td>4.13</td>
<td>The abundance pattern of a typical very metal-poor DLA</td>
</tr>
<tr>
<td>4.14</td>
<td>Comparing the typical abundance pattern of a very metal-poor DLA to models of Population III nucleosynthesis</td>
</tr>
<tr>
<td>5.1</td>
<td>Selected absorption lines from the carbon-enhanced DLA</td>
</tr>
<tr>
<td>5.2</td>
<td>Element abundance pattern in the carbon-enhanced DLAs</td>
</tr>
<tr>
<td>5.3</td>
<td>Ionisation corrections for the carbon-enhanced DLAs</td>
</tr>
<tr>
<td>5.4</td>
<td>An upper limit for the gas temperature of the CEMP DLA</td>
</tr>
<tr>
<td>5.5</td>
<td>A close inspection of the C II line profiles</td>
</tr>
<tr>
<td>5.6</td>
<td>Curve-of-growth for a Doppler parameter of 2.36 km s^{-1}</td>
</tr>
<tr>
<td>5.7</td>
<td>Monte Carlo realisations to test the significance of the carbon-enhancement</td>
</tr>
<tr>
<td>5.8</td>
<td>Comparing the elemental abundance pattern of the CEMP DLA to CEMP stars</td>
</tr>
</tbody>
</table>
5.9 Comparing the abundance pattern for the CEMP DLA to models of Population III nucleosynthesis ........................................ 133

6.1 The explosive energy for carbon-enhanced models of Population III nucleosynthesis. .................................................. 139

6.2 The CEMP fraction as a function of the power law index for the Pop III IMF. ......................................................... 140

6.3 Identifying carbon-enhanced DLAs from the N/Al ratio. .......................................................... 141

6.4 The N/Al ratio from carbon-enhanced models of Pop III nucleosynthesis. ......................................................... 142

6.5 Determining the explosion energy of the first stars. .......................................................... 143

6.6 Determining the amount of mixing in the first stars. .......................................................... 145

7.1 The spatial distribution of SNe Ia from the Union compilation .......................................................... 149

7.2 Deviations of the SNe Ia from the Hubble law. .......................................................... 151

7.3 Model fits to the dipolar model .................................................................................................................. 156

7.4 Results from the maximum likelihood analysis as a function of the angle from the preferred direction of acceleration .................................................................................................................. 157

7.5 Distribution of distance modulus errors in SNe Ia data .......................................................... 158
List of Tables

3.1 Derived e-folding scale lengths of DLAs ........................................ 49
3.2 Metal lines detected at the redshift of the DLA in UM 673A .............. 56
3.3 Absorption components of low ion transitions in the DLA in line to UM 673A .................................................. 58
3.4 Ion column densities of the DLA in UM 673A .................................. 58
3.5 Absorption components of high ion transitions near the DLA in UM 673A ............................................................. 61
3.6 High ion column densities at redshifts close to that of the DLA in UM 673A ............................................................. 61
3.7 Element abundances in the DLA towards UM 673A ......................... 62

4.1 Journal of observations for the very metal-poor DLA survey ............. 75
4.2 Absorption components of low ion transitions ............................... 77
4.3 The adopted metal line laboratory wavelengths and oscillator strengths . 78
4.4 Ion column densities of the DLA in J0311−1722 at $z_{\text{abs}} = 3.73400$ . 81
4.5 Ion column densities of the DLA in J0831+3358 at $z_{\text{abs}} = 2.30364$ . 82
4.6 Ion column densities of the DLA in J1001+0343 at $z_{\text{abs}} = 3.07841$ . 84
4.7 Ion column densities of the DLA in J1037+0139 at $z_{\text{abs}} = 2.70487$ . 86
4.8 Ion column densities of the DLA in J1340+1106 at $z_{\text{abs}} = 2.50792$ . 88
4.9 Ion column densities of the DLA in J1340+1106 at $z_{\text{abs}} = 2.79583$ . 91
4.10 Ion column densities of the DLA in J1419+0829 at $z_{\text{abs}} = 3.04973$ . 91
4.11 C, N, O, Al, Si, and Fe Abundance Measurements in VMP DLAs ....... 94
4.12 The mean and dispersion in X/O for the elements in the VMP DLAs survey .................................................. 104
4.13 Column densities estimated from high and medium spectral resolution spectra ............................................ 109

5.1 Metal lines in the $z_{\text{abs}} = 2.3401$ DLA towards J0035−0918 ............ 118
5.2 Ion column densities in the $z = 2.3401$ DLA towards J0035−0918 ....... 121
5.3 Element abundances in the $z = 2.3401$ DLA towards J0035−0918 ...... 121

7.1 Parameter fitting results ................................................................. 155

A.1 Adopted solar abundances ............................................................ 163

B.1 C, N, O, Al, Si, and Fe column densities in VMP DLAs .................. 166

C.1 [Fe/H] and [O/Fe] in VMP DLAs .................................................. 167
C.2 [Fe/H] and [O/Fe] in metal-poor stars .......................................... 168
Before the Universe was filled with the rich diversity of galaxies, stars, and planets that we see today, it was a cold, dark and rather boring place. The Universe’s baryonic content consisted of nothing more than hydrogen, helium and their isotopes, together with trace amounts of lithium. The more familiar elements on the periodic table, including all elements heavier than helium (which astronomers collectively term ‘the metals’), were nowhere to be seen. Then, the first stars were born, making the Universe a hotter, brighter, and rather more interesting place. This first generation of stars were the first bound structures in the Universe, and created the photons that mark the final frontier in observational astronomy. These photons commenced the epoch of reionization, causing the Universe to undergo the second of its major phase changes. Furthermore, this generation of stars played a fundamental role in creating the complex Universe we see today; at the end of their lives, these stars polluted the Universe with its first dose of metals, allowing low mass stars, like our Sun, to form.

1.1 The birth, life, and death of the first stars

Despite the key role the first stars played in shaping our Universe, very little is known about them. In fact, their existence is solely based on theoretical prediction and indirect observational evidence — no metal-free star has been directly observed to date. Their elusive nature suggests that either: (1) no metal-free stars have ever existed, (2) metal-free stars existed, but have long-since died out, or (3) metal-free stars exist today but we are looking in all the wrong places!
Within the current cosmogonical framework, the first of these possibilities is almost certainly ruled out. The remaining two possibilities depend solely on whether or not metal-free stars exist today, which in turn depends on the time of their birth and how long they lived (or equivalently their birth mass).

Indeed, it seems unlikely that pockets of absolutely pristine, primordial gas still exist in the present Universe; such pristine pockets would need to arise in an ionised medium well before the reionization of the Universe to suppress star formation until later times. Thus, these pockets would need to reside in the \( \text{H}^\text{II} \) region surrounding a nearby, massive star-forming galaxy at early times. In this event, the pristine gas would have likely succumbed to metal enrichment through the winds of their star-forming host galaxy prior to a redshift \( z \sim 2 \) (Johnson, 2010). Metal-free stars (commonly referred to as Population III or Pop III stars for short), therefore, will likely not form today.

If, however, a Pop III star was to form at \( z \approx 2 \) (roughly 10 billion years ago), it would need to be a low mass star (\( \lesssim 1.0 \, M_\odot \)) in order to have survived until the present day. Such a scenario is difficult to envisage, since all evidence to date suggests that metal-free stars are significantly more massive, owing to the absence of cooling by metal-line emission (see e.g. Bromm & Larson 2004). It therefore seems that all metal-free stars in our Universe have been born, lived their lives, and long since died; studying their properties directly with observational techniques could therefore prove difficult, if not impossible.

Perhaps the best handle on the formation and evolution of the first generation of stars has been afforded through computer simulations. Unlike Population I star formation in the local well-observed Universe, which is a highly complex physical process (see e.g. the recent review by McKee & Ostriker 2007), understanding Population III star formation is a relatively straightforward task. One essentially requires two ingredients: (1) an understanding of structure formation, and (2) an accounting of all the relevant atomic and molecular physics in the early Universe. The latter is quite simple; the early Universe contained nothing more than the primordial composition of H, He and their isotopes. Since there were no metals nor dust grains at this time, the pristine gas first cooled primarily through the rotational lines of molecular hydrogen.

The remaining ingredient required to model the birth of the first stars (i.e. structure formation) is modelled in computer simulations with initial conditions set by the observations of the cosmic microwave background (CMB) radiation, and the distance-redshift relation of distant type Ia supernovae (Komatsu et al., 2011; Riess et al., 1998; Perlmutter et al., 1999). These observations suggest that our Universe is well-described by a \( \Lambda \text{CDM} \) cosmology, where most of the matter is comprised of Cold Dark Matter (CDM), but is essentially dominated by dark energy (\( \Lambda \)), possibly in the form of Einstein’s cosmological constant.

Despite these simplistic and well-established physical principles, it has not yet been possible to model the formation of a fully formed Pop III star from cosmological initial conditions. Nevertheless, several authors in the last decade have obtained a valuable insight into primordial star formation, by following the evolution of a collapsing cloud of primordial gas (Bromm,
1.1 The birth, life, and death of the first stars

Figure 1.1: A temperature-density phase diagram from the final simulation output for a collapsing cloud of primordial gas. The gas is first adiabatically heated to the virial temperature of the halo ($T_{\text{vir}}$, shown by the long-dashed horizontal line), at which point the gas cools by molecular hydrogen to a characteristic temperature and density (the bulge of points at $T_c = 200$ K, $n_c \sim 10^3$ cm$^{-3}$). Reproduced from Bromm, Coppi, & Larson (1999).

Coppi, & Larson, 1999; Abel, Bryan, & Norman, 2000). These studies agree that the cloud first undergoes a stage of adiabatic contraction, which heats the pristine gas to the virial temperature of the halo (see Figure 1.1). The gas density has now risen to $n \sim 1$ cm$^{-3}$, which increases the production of molecular hydrogen – the dominant coolant of primordial gas. The lowest energy radiative transition of molecular hydrogen, and therefore the lowest temperature that the gas can eventually cool to, is the transition from $J = 0$ to $J = 2$ which corresponds to a temperature of $E/k_B = 510$ K. The gas thus cools thereafter via molecular hydrogen to a temperature of $\sim 200$ K. This temperature is somewhat less than the lowest energy transition allowed by H$_2$, since the random motions of some particles are drawn from the high energy tail of the Maxwell-Boltzmann velocity distribution.

The gas has now reached a ‘characteristic state’, corresponding to a temperature of $T_c = 200$ K and a density of $n_c \sim 10^3$ cm$^{-3}$. At even higher densities, cooling would become inefficient due to collisional de-excitations; the gas thus loiters in this characteristic state, as indicated by the bulge in Figure 1.1. Eventually, this pristine cloud of gas accumulates enough mass to become gravitationally unstable, which occurs once the mass of the cloud reaches the Jeans mass. Beyond this point, the gas attains higher densities, which further promotes the production of molecular hydrogen and therefore enhances the cooling rate. Finally, a small protostellar core
is born which is the progenitor of the first massive star. This core gradually accretes gas from its surroundings at a rate approximately equal to the Jeans mass divided by the freefall time \( \dot{M} \sim M_J/t_{ff} \approx c_s^3/G \), which is proportional to \( T^{3/2} \). Therefore, the limited cooling afforded by metal-free (as well as dust-free) gas implies that the first protostellar cores had larger accretion rates than today’s more metal-rich stars and hence resulted in larger final stellar masses. It remains to be seen at what stage this accretion phase is terminated, either by radiative feedback or by lack of consumable gas. Nevertheless, most studies agree that the first stars are expected to be more massive than Population I and II stars, with masses typically in the range \( M = 60 - 320 M_\odot \) (see e.g. McKee & Tan 2008).

Due to the recent advances in simulation power, it has become possible to follow the evolution of these pristine birth clouds, starting from cosmological initial conditions through to the near-final stages of Pop III star formation, at a time when the surrounding gas is being accreted onto the protostar. Such simulations use a hierarchical zoom-in procedure to resolve the gas kinematics down to a resolution limit of 50 AU (see Figure 1.2). This increase in numerical resolution has afforded a rather surprising new view on Population III star formation; following the formation of a small protostellar core, the gas collapses to form a nearly Keplerian disc which becomes gravitationally unstable (see Figure 1.3; Stacy, Greif, & Bromm 2010), and forms a Pop III binary or small multiple (see also Turk, Abel, & O’Shea 2009). Clearly, if Pop III stars form in small multiples, this will have a dramatic impact on the final amount of gas that is accreted by these stars, and hence, their final mass. Not only are these protostars competing with their companions for the gas available for accretion, but they must contend with the radiative feedback effects from nearby, more massive companions which could slow or halt the star’s accretion flow.

Such radiative effects are perhaps the single most important ingredient still missing from the current generation of cosmological first star formation simulations. Not only do these radiative effects limit the final mass of the first stars (i.e. negative feedback), but their presence can also yield positive feedback effects. For example, the resulting H\textsc{ii} regions surrounding the first stars are highly ionised, and thus contain a large number of free electrons – which are catalysts for molecule formation. Furthermore, the now increased H\textsubscript{2} abundance acts to increase the abundance of deuterium hydride (HD), which can allow the gas to cool to the temperature of the CMB (Johnson & Bromm, 2006). The cooler primordial gas can now give birth to metal-free stars of much lower masses (\( \sim 10 M_\odot \)) than those described above.

These different formation channels of metal-free stars thus result in Pop III stars with a much wider distribution of stellar masses; this distribution is commonly referred to as the primordial initial mass function (IMF). In the local relatively metal-rich Universe, the IMF for a generation of stars is typically of the form suggested by Salpeter (1955), where the number of stars (dN) in the mass range \( M \) to \( M + \) dM is \( dN \propto M^{-(1+\gamma)}dM \) (where \( \gamma = 1.35 \) for a Salpeter IMF). Such a simple form of the IMF is perhaps less likely to apply to metal-free stars. Unlike star formation in the local Universe, where a large number of stars are born in a single molecular cloud (hence
Figure 1.2: The final output for a simulation of the first stars (on various physical scales) using a hierarchical zoom-in procedure. The simulation box size in each case is provided at the bottom of each panel. The asterisk in the middle and bottom panels denotes the most massive sink particle in the simulations, which is taken to be the protostellar seed for the first star. In the bottom right panel, the second most massive sink particle is represented by the ‘+’ sign. Other sink particles are shown as empty diamonds. Reproduced from Stacy, Greif, & Bromm (2010).
Figure 1.3: An edge-on (left panel) and face-on (right panel) view for a simulation of the metal-free stars, captured 5000 yr after the formation of a protostellar core. The arrows indicate the direction and magnitude of the velocity field. Symbols have the same meaning as the bottom-right panel of Figure 1.2. Reproduced from Stacy, Greif, & Bromm (2010).

resulting in a well-sampled IMF), very few stars are borne out of metal-free gas within a single halo. Therefore, the ‘cosmic’ primordial IMF is likely different from the primordial IMF of each individual halo that forms Pop III stars. Indeed, this suggests that one’s definition of the primordial IMF will depend on the application. For example, when analysing the properties of the metal enrichment of the early intergalactic medium (IGM), one might consider the ‘cosmic’ primordial IMF (see e.g. Madau, Ferrara, & Rees 2001), whereas a study that focuses on the enrichment of a single halo might consider, say, a Gaussian-like IMF centred on the stellar mass that dominates the local metal enrichment (see e.g. Bland-Hawthorn et al. 2011). Regardless of the environment or the functional form of the IMF, however, there is a good general agreement between different studies that suggests the metal enrichment by Population III stars is dominated by massive stars ($\sim 10 M_\odot$; Karlsson, Bromm, & Bland-Hawthorn 2011); the IMF for Population III stars is expected to be ‘top-heavy’ when compared with today’s Population I and II stars.

Massive stars have the ability to synthesise metals on a relatively short timescale ($\sim 30$ Myr), and rapidly deliver these nucleosynthetic products to the surrounding pristine gas through energetic explosions at the end of their lives. The final fates for Population III stars of different mass are illustrated in Figure 1.4, where I show the initial-final mass function of metal-free stars. Below $\sim 9 M_\odot$, a metal-free star evolves in a similar fashion to their more metal-rich counterparts (e.g. Siess, Livio, & Lattanzio 2002); most of their yield comes from winds when the star is on the asymptotic giant branch. The star later dies as a white dwarf. Just above this mass cutoff, there is a small window between roughly $9 - 10 M_\odot$, where a degenerate O/Ne/Mg core forms (see e.g. Wanajo et al. 2009), and the star either loses its outer layers in a wind and hence forms a white dwarf, or collapses due to electron-capture in the core and explodes.
1.1 The birth, life, and death of the first stars

Figure 1.4: The initial-final mass function for non-rotating metal-free stars is represented by the thick solid black line. The mass therefore lost during the star’s evolution is then the difference between the thick solid grey and solid black lines. Reproduced from Heger & Woosley (2002).

Between $10 \sim 100 \, M_\odot$, the central core collapses and the star dies as a core-collapse (Type-II) supernova (see e.g. Heger & Woosley 2010). If the star’s mass is $\sim 25 \, M_\odot$, the core will form a neutron star as a result of the explosion, otherwise the core will form a black hole either directly (for $\sim 40 \, M_\odot$) or by material that is ejected and falls backs onto the central remnant (for $\sim 40 \, M_\odot$). Above $100 \, M_\odot$, a pair-instability develops in the core of the star (see e.g. Heger & Woosley 2002), which acts to reduce the thermal pressure with the production of an electron-position pair. When the mass of the star is less than $\sim 130 \, M_\odot$, the pair-instability causes the star to eject some material from its outer layers and returns to equilibrium. This process repeats itself until the star ‘quietly’ forms a black hole and ejects no metals. For stars above this mass limit, and less than $\sim 260 \, M_\odot$, the first instance of the pair-instability completely disrupts the star, and ends its life as a pair-instability supernova, leaving no remnant behind. For any star above this mass, the pair-instability still causes the star to collapse, which increases the
nuclear energy of the core. Before this energy is large enough to disrupt the star, however, a photodisintegration instability sets in, which causes the star to collapse directly to a black hole — no metals are ejected.

From Figure 1.4, we see that the total mass yield for metal-free stars is given by the horizontal difference between the thick black line and the linear grey line (where $M_{\text{initial}} = M_{\text{final}}$). This ejected mass contains the metals that first enriched the Universe; it is these first metals that provides us with a unique window to study the nature of this first generation of stars. This is made possible by the fact that stars with a different mass expel different nucleosynthetic products at different times and in different ways. One, however, needs to understand the sometimes complex nucleosynthetic and physical processes that created and distributed these metals.

1.2 The nucleosynthesis of the first metals

The dominant metal-yield from metal-free stars is expected to come from massive stars ($10 - 100 \, M_\odot$) that end their lives as core-collapse supernovae (CCSNe). Whilst there have been several detailed models that follow the nucleosynthesis in these metal-free stars, very little is known about the physics behind the explosion mechanism that determines the final metal-yield. Further complications are introduced due to our incomplete understanding of the processes that mix material between neighbouring stellar layers, both during the star’s life and the explosion phase. In this section, I will discuss the latest models that deal with these issues, and highlight the chemical signatures we expect to uncover in gas that was enriched by the first stars.

The first zero-metallicity nucleosynthesis calculations for stars that end their lives as CCSNe were performed by Woosley & Weaver (1995). These yield models follow the evolution and explosion of 12 metal-free stars in the mass range $11 - 40 \, M_\odot$. To simulate the explosion at the end of a star’s life, these authors used a ‘piston’ to deposit momentum at a specified mass coordinate (or equivalently at a given radius). This location was set to be the radius corresponding to a large discontinuous change in the electron mole number due to electron capture, which was typically found to occur at the edge of the iron core. In addition to selecting a location for the piston, one must also choose how to move it during the explosion. Woosley & Weaver (1995) parameterise this motion by an initial velocity that produces a final kinetic energy at infinity of $1.2 \times 10^{51}$ erg for the ejected material. As a consequence of this explosion, the inner regions of the star contract to form a compact object (a neutron star or a black hole) whilst the outer regions are ejected. Some of the ejected material may not, however, overcome the binding energy of the star, and this material falls back onto the central remnant. Thus, the resulting metal yields are calculated at the end of the simulations by considering the final decay products of all unstable nuclei in the material that has enough energy to leave the star entirely.

The model described above is a relatively simple, but physically well-motivated prescription of a CCSN explosion. Woosley & Weaver (1995), however, only include the explosion
1.2 The nucleosynthesis of the first metals

Figure 1.5: A cartoon of the mixing-fallback model for metal-free stars. **Left Panel:** During the explosion, the black region forms a compact remnant and ejects the material in the outermost layers (represented by the white region). The middle layers of the star (interior to the dashed circle, but exterior to the black region) are entirely mixed. The blue region falls back onto the central remnant, whilst the material in the red region is ejected. The metal-yield is therefore represented by both the white and red regions. **Right Panel:** A perhaps more realistic (2D) representation of the mixing-fallback model that is shown in the left panel. In this case, material is primarily ejected along the ‘rotation’ (or jet) axis of the star, whilst material in the rotation plane falls back onto the compact object. Reproduced from Tominaga et al. (2007).

physics that delivers the metals to the surrounding interstellar medium; they have not included the physics which describes how the stellar layers mix during the explosion. This additional physics has large consequences on the final metal yields, particularly that of $^{56}$Ni (which later decays into $^{56}$Fe). If the stellar layers do not mix during the explosion, very little iron would be ejected, since most of the inner material would fall back onto the central remnant.

Such ‘mixing and fallback’ models were first introduced by Umeda & Nomoto (2003) to explain the abundance pattern of the iron-poor yet carbon-rich star HE 0107 − 5240 in the halo of our Galaxy (Christlieb et al., 2002). The basic principles behind these parametric models are illustrated in Figure 1.5. At the end of this example star’s life, a compact remnant is formed out of the material in the black region whilst the material in the white region is completely ejected into interstellar space. The material interior to the dotted line, but exterior to the black region, is assumed to efficiently (and entirely) mix. The red region is expelled, and the blue region falls back onto the central-remnant. Whilst these models are often able to provide an acceptable fit to almost any abundance pattern, they are parametric with little physical motivation; one can essentially find a model made-to-order.

Whilst the physical mechanism behind the ‘fallback’ (i.e. the material that does not have enough energy to escape the binding energy of the star) depends on the unknown explosion mechanism, the ‘mixing’ is primarily thought to be the result of two processes: (1) stellar rotation during the star’s life, and (2) mixing due to the Rayleigh-Taylor (R-T) instability during the
explosion. Whilst both of these effects have received ongoing attention in models of near-solar metallicity CCSNe for the last two decades (i.e. since SN 1987A exploded), only recently have studies emerged that investigate rotational or R-T induced mixing in zero or near-zero metallicity stars. Rotation, for example, is found to have quite startling effects in this regime; not only does rotation provide a means to transport material from the inner to the outer regions of a star, but consequently there are some notable enhancements in the element yields, especially those of CNO (Meynet, Ekström, & Maeder, 2006; Hirschi, 2007). Furthermore, the increased metal abundance of the star’s surface, in combination with the fast rotation rates of low metallicity stars, causes strong stellar winds which can also affect the final metal-yield of zero metallicity stars. Thus, (rapid) stellar rotation in metal-free stars can affect both their physical and chemical evolution in a variety of ways, which in turn determine the final metal yield.

The R-T instability, on the other hand, operates only at the end of a star’s life (i.e. during the SN explosion) and therefore principally affects the metal yield through the mixing it promotes. If the forward shock due to the explosion encounters a region where the density is shallower than $r^{-3}$, it decelerates and thus causes material to build-up behind the shock. This reverses the pressure gradient and a reverse shock forms, which propagates towards the centre of the star (in the frame of the forward shock). The R-T instability forms in this post-shock medium, where the pressure gradient opposes that of the density, and the material develops R-T ‘fingers’ (see Figure 1.6). The material remains unstable, and hence promotes R-T mixing, until the reverse shock passes.

The mixing induced by the R-T instability in metal-free stars is less complete than in stars of solar-metallicity (Joggerst, Woosley, & Heger, 2009), since the former are typically more compact and thus the instability has less time to develop. This does not of course imply that the R-T instability has less importance in mixing the stellar layers. In fact, incorporating both forms of mixing in models of metal-free nucleosynthesis has only been achieved relatively recently (Joggerst et al., 2010a). For the different rotation rates considered by these authors, they found that the final structure of the metal-free star remains relatively unchanged, implying that rotational mixing is perhaps less important. Introducing even a modest rotation rate, however, enlarges the ‘compact’ zero-metallicity progenitors. Consequently, the reverse shock during the explosion now takes a longer time to propagate through the ‘larger’ star. This in turn means that the R-T instability has a longer time to develop, and thus by introducing rotation, the star undergoes more mixing via the R-T instability.

The above simulations are computationally very expensive, since one needs to follow the evolution in two spatial dimensions. Thus, there are no detailed studies to date that consider all of the above effects for a large number of stars that cover the mass range of CCSNe. In order to obtain a better handle on the nucleosynthetic signature of the entire first generation of stars, one requires a high mass resolution study to appropriately sample the IMF. Such a study was conducted recently by Heger & Woosley (2010), who simulated the evolution of metal-free stars of mass $10 - 100 M_\odot$ in steps of $\sim 0.1 M_\odot$. These authors simulate the CCSN explosion
1.2 The nucleosynthesis of the first metals

Figure 1.6: A simulation of Rayleigh-Taylor mixing during the explosion of a \(25 \text{ M}_\odot\) metal-free star. This snapshot shows the logarithm of the mass density at a time when all material is stable (i.e. after the reverse shock has passed). Note that the Fe core (red) has been efficiently mixed into the outer regions. Reproduced from Joggerst, Woosley, & Heger (2009).

Using the piston method adopted by Woosley & Weaver (1995) (described above), located at either near the outer edge of the iron core (as before) or at the discontinuity in density near the base of the oxygen burning core. Mixing in these models is (perhaps unphysically) modelled by a running boxcar filter, which for the ‘standard’ case adopted by Heger & Woosley (2010) has a width of 10% the mass of the He core. This standard model also assumes that all metal-free stars provide a kinetic energy at infinity of \(1.2 \times 10^{51}\) erg.

In the top panel of Figure 1.7, I present the element yields from this standard case, integrated over a Salpeter-like power-law IMF (see Section 1.1) for a range of spectral indices. Specifically, I plot the elemental abundance (as opposed to the yield) as a function of the atomic
Figure 1.7: The abundance for each element, X, relative to Fe is shown for core-collapse supernovae (top panel) and pair-instability supernovae (bottom panel). These yields are integrated over a Salpeter-like IMF with index $\gamma = 0.35$ (red stars), $\gamma = 1.35$ (Salpeter; green stars), and $\gamma = 2.35$ (blue stars). The dashed line corresponds to the adopted solar abundance scale (see Appendix A).
number\(^1\). A few notable deviations from the solar scale (dashed line) are apparent. First and foremost, the abundance pattern exhibits a clear odd-even effect (see e.g. Heger & Woosley 2002), where the odd atomic number elements are deficient relative to the even atomic number elements. This effect is due to the absence of a neutron source, which is required to synthesise the elements that are not produced by \(\alpha\)-capture. Furthermore, for the range of IMFs considered, this figure suggests that the \(\alpha\)-capture elements are enhanced relative to the Fe-peak elements, which is more apparent for shallower IMFs. A related observation is that a shallower IMF (i.e. including more massive stars) results in the production of more elements with an atomic number below \(\sim 15\).

Quite a different view is offered by the more massive metal-free stars that ended their lives as pair-instability supernovae (PISNe). The yields for these stars (from Heger & Woosley 2002), integrated over a Salpeter-like IMF as before, are presented in the lower panel of Figure 1.7. The abundance pattern exhibited by this range of masses provides a different nucleosynthetic signature. First, the odd-even effect is now much more distinct and the lighter \(\alpha\)-elements are not as enhanced relative to the Fe-peak elements. In fact, below an atomic number of \(\sim 15\), the ratio of two elements is virtually independent of the choice of IMF. However, in contrast to the abundance patterns for a population of CCSNe, a shallower IMF for this range of masses would result in fewer elements with an atomic number below \(\sim 20\). Aside from these nucleosynthetic signatures, one of the defining traits of a PISN is the marked enhancement of Si relative to O. All of these signatures of metal-free stars offer a pathway to indirectly study the nature of this now extinct first generation.

### 1.3 The transition to Population II star formation

The ejection of metals and possibly dust (Todini & Ferrara, 2001; Schneider, Ferrara, & Salvaterra, 2004) into the IGM following these first episodes of stellar nucleosynthesis afforded new pathways for the gas to cool below the temperature floor permitted by H\(_2\) or HD cooling in the originally pristine gas. The additional cooling that these pollutants provided enabled the metal-poor gas to condense and fragment to even smaller mass scales, which may have allowed the first low mass, Population II stars to form. Of particular interest is the critical metallicity, \(Z_{\text{crit}}\), of the enriched medium that drove this transition towards the low-mass dominated star formation that we observe in today’s Universe.

At present, there are two competing theories for the dominant physical mechanism that drove this transition: (1) cooling by atomic fine-structure lines which predicts low mass stars to form once the metallicity reaches \(Z_{\text{crit}}/Z_\odot \sim 10^{-3.5}\) (Bromm & Loeb, 2003); and (2) dust-induced fragmentation which instead suggests \(Z_{\text{crit}}/Z_\odot \sim 10^{-6}\) (Schneider et al., 2006). The

\(^1\)Throughout this thesis, I adopt the standard notation: \([A/B] \equiv \log(N_A/N_B) - \log(N_A/N_B)_\odot\), where \(N_{A,B}\) refers to the number of atoms of elements A, B. See Section 2.3 for further details.
Chapter 1. The First Stars

Figure 1.8: The transition discriminant for low mass star formation (dashed line, where the uncertainty in this relation is shown by the dotted lines) is plotted with that measured for stars in our Galaxy (open circles), whilst the red arrows indicate lower limits. The solid line indicates the $D_{\text{trans}}$ corresponding to a solar scaled abundance (see Eq. 1.1). Note that there are no stars residing in the forbidden zone. Reproduced from Frebel & Norris (2011).

dust-induced fragmentation model predicts that gas with just a very mild metal enrichment can form stars down to 0.01 M$_\odot$. Such metal-deficient stars could still be observed today, hence preserving the chemical signature in their stellar atmosphere from the very first episodes of star formation. No such star, however, has yet been observed with such low levels of enrichment. The fine-structure line cooling model by Bromm & Loeb (2003), on the other hand, predicts that gas with carbon and oxygen abundances below $[\text{C/H}] \sim -3.5$ and $[\text{O/H}] \sim -3.0$ is unable to sufficiently cool to form low mass stars. These authors, however, assumed that the fine-structure lines of C$\text{II}$ and O$\text{I}$ act as the dominant coolants, more recent studies (Santoro & Shull, 2006) have shown that including additional cooling due to the fine-structure lines of Si$\text{II}$ and Fe$\text{II}$ introduces only small corrections.

The theory behind the fine-structure line cooling model proposed by Bromm & Loeb (2003) is quite straightforward, and states: the gas will fragment into smaller mass stars if the cooling rate of the gas due to C$\text{II}$ and O$\text{I}$ fine-structure lines exceeds the adiabatic compressional heating rate of a gas at the characteristic density and temperature of pristine gas (see Section 1.1 and Figure 1.1). By equating the balance between heating and cooling, one can formulate a ‘transition discriminant’ (Frebel, Johnson, & Bromm, 2007),

$$D_{\text{trans}} = \log_{10} \left( 10^{[\text{C/H}]} + 0.3 \times 10^{[\text{O/H}]} \right),$$

(1.1)

whereby low mass star formation can proceed in clouds of gas that have a $D_{\text{trans}}$ above the critical value of $D_{\text{trans, crit}} \approx -3.5 \pm 0.2$. Indeed, for all stars that have been observed to date in our Galaxy, this relation seems to hold (see Figure 1.8); no star is observed in the forbidden zone. This is also true for the two most Fe-poor stars known with $[\text{Fe/H}] < -5.0$, which both
1.4 Finding the first metals: Chemical signatures from the Galactic halo

I have to this point painted a picture whereby the first stars are born, synthesise the first metals in their interiors, and later eject these metals into space. These ‘first metals’ are later incorporated into the first Population II stars which might still be alive today if they were born with sufficiently low mass. We are therefore able to build on our understanding of the first stars by searching for the most metal-poor stars in the halo of our Galaxy that still retain the signature from Population III nucleosynthesis.

Perhaps the most obvious place to search for these chemical signatures is the stellar atmospheres of the most Fe-poor stars known, including: (1) HE 0107−5240 (Christlieb et al., 2002), (2) HE 1327−2326 (Frebel et al., 2005), and (3) HE 0557−4840 (Norris et al., 2007). Indeed, the abundance patterns for these three stars show a marked carbon- and oxygen-enhancement relative to iron, all having \([\text{C,O}/\text{Fe}] > +1.0\) (see Figure 1.9). This is perhaps not surprising, since their progenitor cloud must have had these high levels of C and O in order to bear stars with low enough mass to survive until today. This raises an important note of caution: For these stars, one cannot simply integrate the yields of Population III nucleosynthesis over an IMF (see e.g. Figure 1.7) and suitably adjust either the slope or the lower/upper mass cutoffs of the IMF until the best fit is obtained — this will not provide useful information on the universal properties of the first generation of stars (i.e. the IMF for all metal-free stars that formed in the Universe). Rather, it will provide an insight into the mass range of the few stars that polluted the surrounding gas with sufficient C and O so as to later form a low-mass generation of Pop II stars. Indeed, when comparing the abundance patterns of these stars to the model yields of metal-free stars that end their lives as CCSNe, in all cases a Gaussian-like IMF centred on \(\sim 20\)
Figure 1.9: The abundance patterns for the most Fe-poor stars in the halo of our Galaxy are presented (black dots with error bars) and are compared to model yield predictions for the first stars (red stars connected by lines). Top Left: HE 0107–5240, Top Right: HE 0557–4840, Bottom Left: HE 1327–2326, Bottom Right: A typical star with [Mg/H] ≤ −2.9 (see Cayrel et al. 2004). In all cases, the dashed line represents the solar level, and symbols with arrows indicate upper limits. All panels plot [X/Fe] for each element, X, except for the bottom right panel which shows [X/Mg].
1.4 Finding the first metals: Chemical signatures from the Galactic halo

Figure 1.10: The C/O ratio is plotted for disc and halo stars (black circles and blue squares respectively) as a function of their oxygen abundance. Also shown is a Galactic chemical evolution model that considers only the products of Pop II nucleosynthesis (green lines). The solar abundance levels are shown by the dashed lines. Reproduced from Akerman et al. (2004).

M\(_{\odot}\) is favoured with a relatively low explosion energy (6 \(\times\) \(10^{50}\) erg). In hindsight, such a model makes perfect sense; the low explosion energy is able to remove all the outer layers of the Pop III star, but cannot expel a substantial amount of iron from the core. Thus, the fact that a star is Fe-poor does not imply that it is the best representative to probe the Universal nature of the first generation of stars.

Perhaps the best way to uncover the chemical enrichment signature of this (entire) generation, is to consider the limiting values for the abundance ratio of two elements at the lowest metallicity. If such ratios flatten, then this may indicate the global level of enrichment set by the first stellar generation. The first and only study to date that uses this method was performed by Cayrel et al. (2004), who used a sample of 35 halo stars with −4.1 < [Fe/H] < −2.0 to infer the intrinsic nucleosynthetic ratio of 13 elements from the 14 lowest metallicity stars. The abundance pattern from this averaged abundance pattern is shown in Figure 1.9, where I have overplotted the best-fitting model from the suite of CCSN models considered by Heger & Woosley (2010). These results suggest that metal-free stars in the mass range 10 − 15 M\(_{\odot}\) (although favouring 15 M\(_{\odot}\) stars) exploding with 1.2 \(\times\) \(10^{51}\) erg, dominate the first metal enrichment episodes. Such conclusions, however, depend on the uncertain physics that is used to model the CCSN explosion, and are therefore to be taken with a pinch of salt.

One can similarly appeal to the trend in the ratio of two elements at the lowest metallicity to infer the nucleosynthetic products from the first stars; one such example that has been used recently is the C/O ratio. The trend of this ratio in disc and halo stars is illustrated in Figure 1.10. Above [O/H] = −1, C/O rises to near solar values owing to the increased, metallicity dependent carbon yields of Wolf-Rayet stars. This trend is also predicted by models of Galactic chemical
evolution that consider only the products of Population II nucleosynthesis, as shown by the green line in Figure 1.10 (Akerman et al., 2004). When [O/H] < −1, these models predict that C/O should continue to decline, or perhaps plateau, towards the lowest metallicities. Such a trend is not supported by the observations, however, which exhibit a distinct rise in C/O in this increasingly metal-poor regime. There are two competing theories that can explain this upturn at the lowest metallicities: We are probing either (1) the residual carbon-enhanced signature from the first generation of stars (Akerman et al., 2004), or (2) the enhanced carbon yield from rapidly-rotating, low-metallicity Population II stars (Chiappini et al., 2006). At present, we are unable to distinguish between these two models, and we may need to appeal to other diagnostic ratios before any conclusions can be drawn.

It is worthwhile noting, however, that concerns were first raised about the nature of this trend due to the uncertainties of the measured C and O abundances. Indeed, there are several factors that complicate the task of measuring elemental abundances from the stellar atmospheres of stars — particularly in this metallicity regime (Asplund, 2005). A recent careful accounting of such effects, however, has shown that the increasing trend of C/O at the lowest metallicities is robust (Fabbian et al., 2009a).

These concerns bring into question the suitability of using the most metal-poor stars in the Galactic halo to uncover the first metals. How confident are we that the measured abundances in a star’s atmosphere reflect the abundances of the cloud of gas from which the star condensed? These metal-poor halo stars may have formed and been polluted later via mass transfer from a binary companion or even self-polluted by their own nucleosynthesis. Perhaps a more important point is that we are interested in understanding the properties of the first generation of stars; as we have no indication of when or where the extremely metal-poor halo stars formed, their abundances may be due to the incorporation of the nucleosynthetic products from just a few individual stars, rather than a whole generation.

In the following Chapter, I discuss the objects and the techniques that I will use to circumvent the complications associated with measuring the chemical signature of the first stars from the atmospheres of the most metal-poor halo stars in our Galaxy.
Presumably, there must be a ‘missing link’ between the explosions of the first stars and their products’ later incorporation into the most metal-poor stars in the halo of our Galaxy — a gas reservoir where these metals reside before condensing to form a second generation of stars. Such a system should be cold, quiescent, and self-shielded from the ionising background, since the gas must be atomic before it can form molecular clouds and hence stars. The gas reservoir must also be extremely metal-poor and have undergone little to no in-situ star formation. These properties are satisfied by the most metal-poor, high redshift damped Lyman-α systems.

2.1 Damped Lyman-α systems

In a nutshell, damped Lyman-α systems (DLAs) are reservoirs of neutral gas, defined to have a neutral hydrogen column density, $N(\text{H} \, \text{I})$, in excess of $2 \times 10^{20}$ H I atoms cm$^{-2}$. DLAs are usually observed in absorption against a bright background object; in the case of our Galaxy, the background object is a star, whereas extragalactic DLAs are usually observed against a more distant quasar (QSO). An example of the latter technique is illustrated in the top panel of Figure 2.1. With such large H I column densities, the Lyα absorption line is strongly broadened by radiation damping – hence the name, damped Lyα systems. An example of the Lyα absorption profile for a typical DLA is shown in Figure 2.2. The residing metal ions that are dominant in these systems also produce absorption lines, which allow one to establish the overall gas-phase metallicity.
Figure 2.1: Light that is emitted by a quasar travels through the Universe and is absorbed by the neutral hydrogen gas along the way, giving rise to the ‘noisy’ looking spectrum blueward of the quasar’s Lyα emission line (at $\sim 4900\,\text{Å}$). When the quasar’s light intersects a large reservoir of neutral gas (i.e. a DLA, indicated by ‘intervening gas’), a relatively large portion of the spectrum is absorbed, and is easily identified. The dominant metal ions in the DLA also absorb the quasar’s light, giving rise to several metal absorption lines, some of which can be seen in the spectrum redward of the QSO’s Lyα emission line. Image: courtesy of Michael Murphy.

These absorbers were first targeted by Wolfe et al. (1986), whose goal was to detect the high redshift analogues of today’s spiral galaxies. The original driver for this search came from the results of Bosma (1981), who observed a gradual decrease in the neutral hydrogen column density towards the outer edges of nearby spiral galaxies, with a minimum of $N(\text{H}^1) = 1.8 \times 10^{20} \, \text{cm}^{-2}$ observed near the outer edges. Wolfe et al. (1986) adopted this lower limit in their pioneering study to search for the neutral gas discs of high redshift spiral galaxies. However, as cautioned by Wolfe et al. (1986), it is difficult to distinguish a neutral gas disc from a protogalaxy, since the latter also contains a large column of H1 gas before eventually collapsing to form stars.

Figure 2.2: A zoomed in version of an example damped Lyα line for a DLA (at $\lambda \approx 1215.7\,\text{Å}$). The QSO continuum has been normalised here, and therefore corresponds to a residual intensity of 1.0 (long-dashed line). There are also a number of unrelated weaker absorption features in the wings of the DLA’s Lyα line, which are due to small overdensities of H1 gas along the line-of-sight to the QSO.
Since this initial investigation, many astrophysical fields have benefited from the study of damped Lyman-\(\alpha\) absorption systems. As reviewed by Wolfe et al. (2005), DLAs have provided key clues into the formation of galaxies at high redshift, clues that are unobtainable by any other means. For example, such systems are found to contain the majority of neutral gas in the Universe, thus suggesting that DLAs are the reservoirs from which stars will later form. Furthermore, DLAs provide key information on the properties of the residing neutral and molecular gas, as well as the chemical and kinematical properties of high redshift galaxies with low star formation rates. In the remainder of this Chapter, I will review the current state of DLA research that is relevant to the work presented in this thesis.

### 2.2 Host Galaxy Identification

We are still unable to say, with confidence, whether a given DLA is probing the extended HI disc of a galaxy, a smaller subgalactic size halo, or a small cloud of HI gas within a much larger galaxy. Perhaps the most direct way to distinguish between these possibilities would be to observe in emission the host galaxy that gives rise to the DLA. Unfortunately, this is very challenging; whilst DLAs are somewhat straightforward to identify spectroscopically, they have proved to be extremely elusive to observe in emission. In fact, the morphologies of very few absorbers at high redshifts (\(z_{\text{abs}}>1.0\)) have been exposed (see Fynbo et al. 2010 and references therein). This difficulty stems from the need to subtract the light from a bright, background QSO in order to reveal the relatively faint emission from the DLA’s host (an example is presented in Figure 2.3). This is only possible, however, when the host galaxy is sufficiently bright.

One could alternatively search for the hosts of low-redshift DLAs with the hope that such galaxies are similar to those that host DLAs at high redshift, some 10 billion years ago. This has become possible recently with the advent of the Cosmic Origins Spectrograph on the Hubble Space telescope, which is sensitive down to an observed wavelength of 1150 Å. This instrument was used by Meiring et al. (2011) in their survey for low redshift DLAs along the line of sight to 37 QSOs. These authors found just 3 DLAs along these sightlines; follow-up imaging of this field revealed several possible candidates, including a large spiral galaxy centred on the QSO (see Figure 2.3). Subsequent spectroscopy of the candidates in the field revealed the spiral galaxy as the host of the QSO, not the DLA. Furthermore, none of the other candidates were found to have a similar redshift to the DLA. Thus host identification is difficult, even at low redshift! Certainly, the lack of DLA host galaxies that have been identified suggests that the cross-section for DLAs is largely contributed by low luminosity galaxies.

Greater (but still limited) success in determining the identity of the host galaxy has been afforded through integral field spectroscopic studies of the field along the line of sight to the quasar (see e.g. Péroux et al. 2011). By adopting this technique, one is able to cleanly separate the line emission of the host galaxy from the QSO continuum, and thus obtain an independent
Figure 2.3: An image of the field towards the QSO J1009 + 0713. After fitting a point spread function to the quasar, and subtracting the result, a galaxy is revealed lying on top of the quasar with an H II region (indicated by the white box). Spectroscopic follow-up reveals that the galaxy is associated with the quasar and not the DLA. Other galaxies that are labelled in the nearby field also differ in redshift to the DLA. The arrow points to blob, slightly offset from the subtracted QSO, which may be the host galaxy of the DLA, however, this candidate requires spectroscopic confirmation. Reproduced from Meiring et al. (2011).

measure of the redshift to confirm the identification. In many cases, the host is identified in Hα emission, or the forbidden [O II] and [O III] emission lines. In fewer cases, one can observe the emission from the Lyα line in the trough of the damped Lyα absorption profile (Hunstead et al., 1990; Warren & Moller, 1996; Möller, Fynbo, & Fall, 2004). The latter is considerably challenging, however, since Lyα is a resonance line and is efficiently absorbed by dust grains. The large number of H I atoms in DLAs efficiently scatter a Lyα photon until it encounters a grain of dust, at which point it is absorbed and radiated away as infrared radiation.

The observed line emission can in some cases provide a useful diagnostic for the rate at which stars are formed (Kennicutt, 1998). This can in turn be used as a very rough proxy for the rate at which the surrounding gas is enriched with metals, in the sense that higher star formation rates would imply a higher metal content. This simple, approximate relation would then imply that the most metal-deficient systems are (perhaps obviously) those that are forming the least stars, and are much less-likely to be polluted by subsequent generations of star formation. This picture is certainly supported by cosmological simulations of galaxy formation (Pontzen et al., 2008; Tescari et al., 2009), whereby the most metal-poor DLAs have undergone little to no in
2.2 Host Galaxy Identification

Figure 2.4: A cosmological simulation of galaxy formation, which highlights the gas that we would observe as a DLA (shades of red). Reproduced from Pontzen et al. (2008).

In fact, such simulations provide the clearest view into the nature of the galaxies that host DLAs. Recent advances in computational techniques have permitted high mass resolution simulations that incorporate the gas physics relevant to the present discussion. For example, the simulations investigated by Pontzen et al. (2008) use a smooth particle hydrodynamics code to simulate a dwarf galaxy, as well as a Milky Way type spiral galaxy at $z = 0$. These simulations take into account the effects of gas cooling and star formation feedback (the latter includes a prescription for depositing metals in the ISM through supernova explosions). After following the simulations to $z = 0$, Pontzen et al. (2008) then trace the simulations back in time to $z = 3$, a redshift where DLAs are commonly observed with ground-based optical telescopes, and study the properties of DLAs at this epoch.

In Figure 2.4, I show the $z = 3$ snapshot from the simulations of Pontzen et al. (2008), which depicts the system that evolves into a Milky Way type spiral galaxy at $z = 0$. For this example, all sightlines through this system that we would perceive as a DLA are shown by the shades of red. It can be readily seen that the brightest galaxy in this field (i.e. the parent halo at $x = 0$, $y = 0$) is not the only halo that could give rise to DLA absorption. Thus, studies
that use imaging or integral field spectroscopic techniques to identify the ‘host galaxy’, might instead simply identify the brightest, nearby galaxy. Indeed, the impact parameters for the ‘host galaxies’ of DLAs identified to date are typically up to several tens of kpc (see Table 4 from Péroux et al. 2011).

Cosmological simulations have thus provided a useful, complementary insight into the nature of DLAs, with the general conclusion being that DLAs are a ‘mixed bag’ of objects. Even though firm evidence is yet to surface regarding the individual nature of these high column density absorbers, a picture is slowly evolving whereby the only property that the entire DLA population has in common, is their high column density of neutral hydrogen.

2.3 DLAs as Chemical Laboratories

The information available on the chemical properties of DLAs has come from studies that use spectroscopy to measure the absorption lines of the residing chemical elements. In this section I will explain the ‘toolkit’ that is used to study the composition of DLAs, in addition to the advantages of using such systems to probe the chemical makeup of systems over a range in metallicity and redshift.

2.3.1 Absorption line formation

The first advantage of using DLAs is that their physics is simple and the derived abundances do not depend on the shape, size, or mass of the absorber. Moreover, the elemental abundances are straightforward to derive, quite unlike measures of elemental abundances in metal-poor Galactic halo stars which, as mentioned in Section 1.4, have their limitations. The only underlying assumption is that the random motions of the atoms residing in the DLA obey a Maxwell-Boltzmann distribution. Thus, the observed absorption line profiles are described by the convolution of the intrinsic line profile of an atom (which is a Lorentzian) with a Maxwellian profile, characterised by a cross section for absorption

\[ \sigma_v = a_0 H(a, x), \]  

(2.1)

where \( a_0 \) contains the atomic parameters of the transition and \( H(a, x) \) is the Voigt integral, given by:

\[ H(a, x) = \frac{a}{\pi} \int_{-\infty}^{\infty} \frac{\exp(-y^2)}{(x-y)^2 + a^2} dy \]  

(2.2)

\[ a_0 = \frac{\sqrt{\pi} e^2}{m_e c^2 \Delta v_D} \]  

(2.3)
where $f$ is the oscillator strength. The damping parameter of the intrinsic line shape, $a$, and the Doppler frequency, $\Delta \nu_D$, are of the form

$$a = \frac{\Gamma}{4\pi \Delta \nu_D}$$

$$\Delta \nu_D = \frac{b}{\lambda_0} = \frac{1}{\lambda_0} \sqrt{b_{th}^2 + b_{turb}^2},$$

where $\Gamma$ is the transition rate, $\lambda_0 \equiv c/\nu_0$ is the rest wavelength of the transition, $b_{turb}$ is the turbulent Doppler parameter of the gas (describing the random motion of the atoms), and $b_{th}$ is the thermal Doppler parameter (describing the thermal broadening of the line profile) which is given by

$$b_{th} = \sqrt{\frac{2kT}{m_{atom}}}$$

Finally, the dimensionless parameter $x$ in the Voigt integral (Equation 2.2) describes the frequency offset from the line centre, in units of the Doppler frequency,

$$x = \frac{\nu - \nu_0}{\Delta \nu_D}.$$

and $\gamma$ is the convolution parameter. Thus, the optical depth for a given atom in the cloud is given by $\tau_{\nu} = N \sigma_\nu$, where $N$ is the column density of this atom. The flux from the quasar, $I(\nu)_0$, is then diminished by a factor of $e^{-\tau_\nu}$ centred on the redshifted wavelength of the absorption line $\lambda_{\text{obs}} = (1 + z_{\text{abs}}) \lambda_0$, and the observed line profile is then given by

$$I(\nu) = I(\nu)_0 e^{-\tau_\nu}.$$

To illustrate the form of this line profile, in Figure 2.5 I present an example profile from three regimes, $\tau_{\nu_0} \ll 1$, $\tau_{\nu_0} \gtrsim 1$, and $\tau_{\nu_0} \gg 1$.

The above formalism has not yet taken into account the broadening due to the instrumental
Figure 2.6: The curve of growth for the Lyα line is shown for three choices of the Doppler parameter: $b = 3 \text{ km s}^{-1}$ (solid line), $b = 10 \text{ km s}^{-1}$ (dashed line), $b = 30 \text{ km s}^{-1}$ (dotted line). All three lines have been colour-coded depending on whether an absorption line would reside on the linear (red), flat (green), or damped (blue) regimes of the curve of growth (cf. Figure 2.5). Note that there is nearly a direct correspondence between the equivalent width of an absorption line and the column density in the linear and damped regimes.

The observed flux spectrum is then given by $I(\nu)_{\text{obs}} = \Phi(\sigma_{\text{inst}}) \otimes I(\nu)$. In the event that $\sigma_{\text{inst}} \gg b/\sqrt{2}$, the spectral resolution is not sufficient to resolve the intrinsic line profile. In this instance, we can simplify the interpretation of the above line profile, by introducing an equivalent ‘saturated’ (i.e. zero flux everywhere), rectangular line profile, such that the area of the rectangular line profile is equal to the area between the true line profile and the continuum. The width of this line profile, known as the equivalent width ($EW$), is then of the form

$$EW = \int_{0}^{\infty} \frac{I(\nu)_{0} - I(\nu)}{I(\nu)_{0}} \, d\nu = \int_{0}^{\infty} (1 - e^{-\tau_{\nu}}) \, d\nu \quad (2.9)$$

Even though the Doppler parameter, $b$, cannot be measured directly under these conditions, one can relate the $EW$ to the column density, and represent their relation on a ‘curve of growth’ (COG). Figure 2.6, depicts three such models with different Doppler parameters. This figure has several notable features; for optical depths at the line core $\tau_{\nu_{0}} \ll 1$, the absorption feature resides on the linear part of the COG, and the column density that is derived from the measured $EW$ is essentially independent of the Doppler parameter. Increasing the optical depth to $\tau_{\nu_{0}} \gtrsim 1$ saturates the line profile and measurements of $b$ and $N$ are degenerate; this is known as the flat part of the COG. One can only break this degeneracy by either measuring the column density
from the linear regime, or by using several transitions of different strength from the same ion to constrain both $b$ and $N$ simultaneously. Increasing the optical depth further still has little affect on the line width until $\tau_{00} \gg 1$ and the damping parameter dominates the $EW$ for the observed line profiles, moving the transition onto the damped regime on the COG. In fact, it is the damping parameter, in combination with a high column density, that gives rise to the broad Ly$\alpha$ line in DLAs.

This feature of DLAs makes them easy to identify in the spectra of quasars, even at low spectroscopic resolution ($R \sim 1000$). Large scale surveys, such as the Sloan Digital Sky Survey (SDSS), have thus uncovered large numbers of DLAs ($\sim 1000$; Noterdaeme et al. 2009; Prochaska & Wolfe 2009). These surveys have increased the number of known DLAs by over an order of magnitude. In contrast to the strong damped Ly$\alpha$ line, however, the relatively weak and complex absorption line profiles exhibited by the metal-lines, are not resolved with the low resolution SDSS discovery spectra. Moreover, the multi-component structure of these metal lines makes it difficult to perform a COG analysis; one must fit a single ‘effective’ Doppler parameter to the DLA’s metal lines, which is almost never true for these systems.

### 2.3.2 Observational techniques and analysis

In order to perform an accurate chemical abundance study, one is required to reobserve these systems with instruments of high spectral resolution ($R \sim 30000$) to resolve the multi-component structure of the metal-lines. This is not a simple task, since the targeted quasars are typically quite faint ($m_r \sim 18$) and require one night of telescope time on an 8−10 m class telescope to achieve the required signal-to-noise to perform an accurate abundance analysis. In large part, the majority of this work has been carried out in recent years with either the Ultraviolet and Visual Echelle Spectrograph (UVES; Dekker et al. 2000) mounted on the European Southern Observatory’s (ESO) Very Large Telescope facility (VLT) in Chile or the High Resolution Echelle Spectrometer (HIRES; Vogt et al. 1994) at the W. M. Keck Observatory’s twin 10 m telescopes on the summit of Mauna Kea. To achieve this high spectral resolution, these instruments employ a technique called echelle spectroscopy, which I now discuss in further detail.

The basic principles of echelle spectroscopy are illustrated in the schematic diagram shown in Figure 2.7. The light collected by the telescope is focused onto the entrance (or viewing) slit and is then passed through a collimator to produce a parallel beam of light. The parallel beam is directed onto a grating which disperses the light. For echelle spectrographs, the high dispersion is generated by the echelle grating, which contains $\sim 50$ lines per millimetre. For comparison, classical long-slit spectrographs contain gratings with $\sim 1000$ lines per millimetre. When light of wavelength $\lambda$ is incident on a grating that has been rotated to an incident angle $i$, then for a given spectral order $m$ and line spacing $d$ the light will leave the grating at an angle $\theta$ governed
by the equation

\[ m\lambda = d \left( \sin i + \sin \theta \right) \quad (2.10) \]

and the corresponding dispersion is given by

\[ \frac{d\theta}{d\lambda} = \frac{m}{d \cos \theta}. \quad (2.11) \]

Thus, for an echelle grating, where \( d \) is relatively large, high dispersions are achieved by using the grating at high orders numbers (\( m = 60–155 \) for UVES; 35–120 for HIRES) and at large angles of diffraction \( \theta \). One consequence of using an echelle grating, however, is that the large dispersion causes the high orders to overlap with one another. Fortunately, for a given line spacing and angle of incidence, the wavelength range will be different for all orders (see Eq. 2.10); one can thus separate the overlapping orders by introducing an additional dispersing element — either a conventional grating or prism(s) — that disperses the light in a direction perpendicular to the high-dispersion axis. A series of lenses then corrects the light and forms
2.3 DLAs as Chemical Laboratories

Figure 2.8: An arc lamp exposure showcasing the CCD mosaic for the upgraded HIRES detector. The blue-sensitive chip is on the left, whilst the red-sensitive chip is on the right.

A symmetric beam which is focused by the camera mirror onto the CCD detector where the data are recorded. An example of the data format from the HIRES instrument is provided in Figure 2.8. The echelle grating disperses the light in the up-down direction, whilst the cross-disperser separates the orders in the left-right direction.

The images are recorded with a charge-coupled device (CCD) that is applied with a small voltage to enable photon counting. Ideally, this should be a constant offset, but in practice there are small variations between all the pixels that are introduced due to the readout of the CCD. To correct for this effect, one must take a zero second exposure to later subtract from all frames. Several bias frames need to be taken and median-combined with an outlier rejection algorithm to statistical sample the bias level and remove contamination from cosmic rays.

After subtracting the bias level, one must remove the pixel-to-pixel variations on the CCD by taking a well-exposed image of a continuous light source, such as the twilight sky or a quartz halogen lamp. Again, several ‘flat-field’ frames are taken and median combined to produce a master flatfield frame. After normalising the continuous light source, one then divides the object exposures by this master flat.

Having removed the additive and multiplicative signatures on the CCD, the individual echelle orders must be extracted. This is achieved by fitting a low order polynomial to either the light from the object frame itself in the case of a point source (such as a quasar), or by taking a flat-field exposure with a pinhole decker. The latter requires a slitlength which is of order the size of the slitwidth. This provides the locations of the orders on the CCD. The trace frame is therefore used as a guide to extract the object’s spectrum. The extracted spectrum not only contains the
photons from the object, but it also contains background sources of light (dominated by the night sky) which must be removed. This step is achieved by selecting a region on the raw frame near to the object’s spectrum and subtracting this ‘background’ from the object.

The result of the above operations is a reduced object spectrum for each spectral order together with an error spectrum that reflects the uncertainty in the observed flux density (including the error due to counting statistics, sky background, and detector noise). Before combining the data from all spectral orders, one must wavelength calibrate the orders by determining a relationship between pixel location on the detector and wavelength. This information is provided by taking an exposure of an arc lamp, typically a Thorium-Argon lamp (as shown in Figure 2.8) since there exist a large number of lines that are narrower than the instrument resolution. The lamp’s spectrum is extracted at the same location as the object’s spectrum, and lines are identified according to the spectral atlas for the arc lamp. For each order, a polynomial is fit to the wavelength as a function of the pixel location, which is then applied to the object spectrum. The above steps were all carried out with reduction pipelines written by ESO in the case of UVES data reduction (the ESO UVES pipeline), and by Tom Barlow in the case of the HIRES data reduction (MAKEE).

The individual echelle orders were then merged into one continuous spectrum manually (for both UVES and HIRES data) using the software package UVES_POPLER maintained by Michael Murphy. At this point, two additional corrections must be applied to the high spectral resolution data to account for (1) the motion of the Earth relative to the Sun, known as heliocentric correction and (2) mapping the data onto a vacuum wavelength scale. Once these corrections are applied, UVES_POPLER resamples the data to a user-specified binsize, and coadds data from several exposures and spectral orders that overlap. One then has the option to remove further defects, such as spikes (e.g. cosmic rays, bad pixels), poorly traced orders, and ghosts (reflections of the object’s spectrum on a traced order). As a final step, the quasar’s continuum and emission lines are normalised by fitting a polynomial to the regions free from absorption features.

The data are now completely reduced and ready to be prepared for analysis. To this end, I wrote a suite of PYTHON routines designed for this purpose. The first step is to derive the DLA’s H\textsc{i} column density from the Ly\textalpha absorption line whilst simultaneously correcting for possible mistakes in the original continuum fit estimated in UVES_POPLER near the broad Ly\textalpha feature. The associated metal lines are then manually identified with reference to an atomic table that lists the rest frame wavelength, oscillator strength, and transition probabilities ($\lambda_0, f, \Gamma$ respectively; cf. Section 2.3.1). Throughout this work, I adopt the values measured in the laboratory reported by Morton (2003) with updates from Jenkins & Tripp (2006). Upon identification, the metal lines that are detected with $\geq 3\sigma$ significance are extracted in small windows of $\pm 150 \text{ km s}^{-1}$ around the pixel with highest optical depth (i.e. lowest flux). Additional metal lines that are undetected with $3\sigma$ confidence are also extracted to calculate an upper limit to the corresponding ion’s column density. A further fine continuum adjustment is performed to all extractions when necessary by fitting a low order polynomial to the nearby continuum.
The extracted lines are simultaneously fit with Voigt profiles using the VPFIT software written by John Webb (Webb, 1987) and maintained by Bob Carswell. A list of the extracted metal-lines are used as input, together with the wavelength range to be fitted and the instrument resolution of the data at the wavelength of the absorption line (usually measured from the narrow arc lines). After providing a set of starting parameters, VPFIT generates synthetic line profiles based on the equations described in Section 2.3.1 and compares these profiles to the observed ones. The difference between computed and observed profiles is then minimised by adjusting the redshift, Doppler parameter, and/or the column density for a given ion. A new set of synthetic profiles is generated and the process iterates until the user-specified convergence criteria is reached (typically when the $\chi^2$ changes by less than 0.05%).

With the standard assumption that the low ionisation metal ions are kinematically associated with the same gas, the redshift and turbulent velocity dispersion are tied for different metal ions in the same cloud. This is a reasonable assumption, since the Voigt profiles used in modelling the DLA absorption in VPFIT correspond to an isothermal cloud of gas. This implies that the gas is in the same fractional ionisation state (i.e. $N$(C II)/$N$(C), for example, is constant) at all points in the cloud. Furthermore, for all reasonable gas temperatures (generally $< 10^4$ K; see e.g. Wolfire et al. 1995), one can assume that the turbulent velocity component in Eq. 2.5 dominates over the thermal velocity component. The final goal is to plausibly fit the data with the minimum number of absorption components, resulting in estimates of the cloud model (including the redshift and velocity dispersion for all fitted components) and the column density for each ion that gives rise to an absorption line.

### 2.3.3 Studying the chemical properties of DLAs

In principle, the column density of a given metal, $N$(M), should be calculated by measuring the column density for all ionisation states of a given element associated with the $\text{H}^1$ gas (i.e. $N$(M) = $N$(M I) + $N$(M II) + $N$(M III) + ...). In practice, this is impossible. Fortunately, the large $\text{H}^1$ column density of DLAs act to self-shield the gas from external radiation fields and, in particular, the UV background radiation from galaxies and quasars (Haardt & Madau, 2001). This renders most metals into a single, dominant ionisation state, and $N$(M) $\approx N$(M D) where D is the dominant ionisation state – typically a neutral or first ion. The column density of neutral hydrogen, which is required to establish the overall metallicity of the gas, is quite simply (see Figure 2.2) and uniquely (see Figure 2.6) derived from the damping wings of the Ly$\alpha$ absorption line profile. The gas-phase metallicity is then found by taking the ratio of a metal-line column density, $N$(M D), to that of hydrogen, $N$(H I), and referring the result to a ‘baseline’ that allows one to easily compare element ratios of different metallicity. This operation is often given the notation $[\text{M/H}] \equiv \log \left(\frac{N(\text{M D})}{N(\text{H I})}\right) - \log (\text{M/H})_\odot$, where the adopted baseline, $(\text{M/H})_\odot$, 

\[1\] In the analysis presented in this thesis, I assume that the gas is in a cold neutral medium with temperatures of a few hundred degrees Kelvin, unless stated otherwise.
is the solar abundance scale. Throughout this thesis, I will adopt the solar abundance scale recently proposed by Asplund et al. (2009), taking the photospheric or meteoritic abundances, or an average of the two, based on the suggestion by Lodders, Plame, & Gail (2009). The adopted abundance scale is provided in Appendix A.

In addition to DLAs having a simple ionisation structure, the derived elemental abundances from narrow metal lines are typically accurate to within 0.1 dex, and are independent of the temperature and density of the gas. The error in the H\textsc{i} column density is also less than $\sim 0.1$ dex. DLAs are thus the best known objects in the high redshift Universe to accurately measure chemical abundances. Results from the first decade of exploring the chemical abundances of DLAs indicate that these systems are typically metal-poor (Pettini et al., 1997b, 1999), having a gas-phase metallicity $\sim 1/10 Z_\odot$ (top panel of Figure 2.9). This observation, together with the known difficulty of observing the host galaxy of DLAs (see Section 2.2), implies that DLAs are not responsible for the majority of star formation at any redshift.

However, as pointed out initially by Pettini et al. (1990), one complication that might arise when measuring the metallicity is the possibility that DLAs hide some fraction of their metals in dust grains. This raises an important question: Do the metal-rich, actively star-forming, and thus dusty DLAs extinguish the light from the background QSO to the point that the QSO is no longer detectable? If this is the case, we would only observe the less-dusty, metal-poor DLAs, thus introducing a bias into the chemical abundance studies of DLAs.

A common method that is used to quantify the effect of dust is to consider a volatile and a refractory element, such as Zn and Cr, and compare their relative abundance to the solar value. Both Zn and Cr are Fe-peak elements, and therefore share similar nucleosynthetic histories, however, Zn shows only a mild affinity to dust grains, whereas Cr is readily depleted. A combination of their proximate wavelengths and relatively weak lines (placing them almost always on the linear regime of the COG), make the Zn\textsc{ii} doublet and Cr\textsc{ii} triplet perhaps the ideal pair of ‘depletion probes’. The first study to investigate the dust content of a DLA by using this pair of elements was carried out by Pettini et al. (1990), who reported on a DLA that is only mildly dusty compared to the interstellar medium of our Galaxy.

Recent compilations of Zn and Cr measurements taken from optically-selected samples of QSOs, such as those reported by Kulkarni et al. (2005), suggest that DLAs suffer a metallicity dependent depletion effect; DLAs with a metallicity less than $\sim 1/100 Z_\odot$ exhibit very little depletion onto dust grains. To test whether or not there exist a population of metal-rich, dusty DLAs missing from optically-selected samples, Akerman et al. (2005) extended the database of Zn and Cr measurements to include DLAs chosen from a radio-selected sample of quasars (see lower panel of Figure 2.9). These authors conclude that neither sample reddened the QSO more than the other. This conclusion is also supported by studies that find a negligible difference in the continuum slope of quasars with and without DLAs along their line of sight (Murphy & Liske, 2004; Frank & Péroux, 2010). These findings suggest that we are likely not missing a population of high-metallicity, dusty DLAs; DLAs are truly chemically unevolved systems.
Figure 2.9: Abundances derived for DLAs based on optically-selected (green symbols Kulkarni et al., 2005) and radio-selected (black symbols Akerman et al., 2005) samples of quasars is shown as a function of the DLA’s redshift. Triangles indicate upper limits on [Zn/H]. Top Panel: DLA metallicity as a function of the DLA’s redshift. Lower Panel: The degree of dust-depletion in DLAs is shown by the ratio of a refractory to volatile element ratio (in this case, [Cr/Zn]). When the DLA’s true metallicity (as traced by the [Zn/H] ratio) reaches $10^{-2}Z_\odot$, the gas in DLAs exhibits very little depletion onto dust grains. Note that in both panels, the radio- and optically-selected samples trace each other well, implying that a population of dusty DLAs are unlikely to be missing from either sample. Reproduced from Akerman et al. (2005).
Figure 2.10: The observed metallicity distribution of DLAs is shown by the blue symbols and error bars. Overplotted in black is the simulated metallicity distribution of DLAs (cf. Figure 2.4). The dotted, dot-dashed, and dashed lines correspond respectively to the metallicity distributions of simulated DLAs with haloes of decreasing virial mass, as indicated by the accompanying legend. Reproduced from Pontzen et al. (2008).

In summary, the last decade of studying DLAs with high resolution spectrographs has amounted to large samples of $\sim 100$ DLAs with accurate abundance measurements (Prochaska et al., 2007; Noterdaeme et al., 2008). The general consensus from these studies is that DLAs are chemically unevolved systems, which is certainly in line with the conclusions in Section 2.2 above regarding their generally low star formation rates. To illustrate this point, in Figure 2.10 I show the observed metallicity distribution of DLAs (blue symbols, from Prochaska et al. 2007), which demonstrates that the typical DLA has a metallicity of $\sim 1/30$ of solar.

2.4 Finding the first metals: Chemical signatures from DLAs

Of particular interest, are the most metal-poor DLAs, which fall in the left tail of this distribution. Such systems, although relatively rare, are perhaps the most suitable candidates for the proposed ‘missing link’ between the yields of the first stars and the incorporation of their metals into the second generation of stars. However, given the long integration times that are required to observe the background quasars, one challenge is to efficiently identify the handful of DLAs that still harbour the signature of Population III nucleosynthesis; one needs to select not only the DLAs that are yet to convert some of their gas into stars (which might pollute the signature), but also those that have not been polluted by external, star forming galaxies.
Cosmological simulations of galaxy formation suggest that the most metal-poor DLAs arise in low mass haloes located in the outskirts of galaxies (see Figures 2.4 and 2.10), which has also been confirmed observationally through a ‘mass-metallicity’ relation (see Figure 2.11). These simulations suggest that such DLAs have converted very little gas into stars when we observe them at redshifts $z \sim 3$ (Pontzen et al., 2008) – they might indeed still retain the signature from the first few generations of stars.

Indeed, it is a great advantage to catch these pristine systems before they form a new generation of stars, like those in the halo of our Galaxy, since DLAs are much simpler physical systems; their abundance measurements are independent of the geometrical configuration and thermodynamical state of the gas. The only pertinent difficulties for DLA abundance measurements are line saturation, and whether dust hides some fraction of a given element. Fortunately, both of these effects are known to introduce smaller corrections to the abundance measurements of DLAs when the metallicity is $Z < 1/100Z_\odot$, which is the metallicity regime we expect to uncover the enrichment signature of the first stars. In addition to their physical simplicity, DLAs are less likely than stars to have their chemical composition altered; it would be difficult to washout the chemical signature of the first stars imprinted on DLAs via external, nearby, star forming galaxies, since DLAs contain $\gtrsim 10^6 M_\odot$ of material. On the other hand, the stellar atmospheres of stars could quite easily be contaminated by mass transfer from a binary companion, or self-polluted by its own nucleosynthesis.

Despite the above-mentioned simplicity with which one can measure chemical abundances
in the most metal-poor DLAs, their rarity has allowed very few studies to uncover their chemical properties. Indeed, to date, the abundance pattern of just one DLA has been compared to the model yield calculations of metal-free stars. This study, by Erni et al. (2006), reported archival observations of the quasar Q0913+072, which has a metal-poor DLA ([Fe/H] = −2.77) along its line-of-sight at \( z_{\text{abs}} = 2.61843 \). By integrating the model yields of CCSNe and PISNe over a power-law IMF, \( dN/dM \propto M^{-(1+\gamma)} \) (where \( \gamma_{\text{Sal}} = 1.35 \) is a Salpeter IMF), these authors find that this DLA’s abundance pattern is consistent with being enriched by a generation of \( 10 – 50 \) \( M_\odot \) metal-free stars exploding as CCSNe at the end of their lives. However, whilst Erni et al. (2006) state that their derived power-law index for the IMF is steeper than Salpeter, \( \gamma > \gamma_{\text{Sal}} \), they neither provide the derived value of \( \gamma \), nor a plot to demonstrate their best fit. Merely, they provide a comparison between the DLAs abundance pattern and the yields from a 15 \( M_\odot \) metal-free star (see Figure 2.12). Certainly, given that the DLA’s abundance pattern is reasonably well-fit by a single, relatively low mass Population III star, the power-law index must by quite steep indeed. Erni et al. (2006) also considered the yields from PISNe in their study, and commented that such stars must not have contributed significantly to the DLAs enrichment; PISNe exhibit a large odd-even effect, that could not be reconciled with the measured N and Al abundances (see Figure 2.12). Finally, whilst these authors found that the abundance pattern of this DLA is consistent with being enriched by a generation of metal-free stars, they were unable to unambiguously rule out a possible contribution from low metallicity Population II stars. Thus, until abundance measurements from additional elements are obtained for this system, we...
are unable to distinguish between enrichment by Population II or Population III stars.

Such studies will have to wait until the next generation of 30+ m telescopes come online in order to efficiently achieve the high S/N ratio that is required to detect the weak lines of Cr, Ni and Zn, which might act as discriminants between Population II and Population III nucleosynthesis. In the meantime, one can consider several diagnostic ratios to infer the nature of the first stars. For example, as outlined in Section 1.4 and Figure 1.10, the C/O ratio in halo stars is found to increase with decreasing metallicity, which could be the residual signature of a high carbon-producing generation of early stars. This picture is corroborated by observations of the most metal-poor DLAs, where the present sample of 6 DLAs all exhibit a C/O ratio inline with the Galactic halo stars (Pettini et al. 2008a; see Figure 2.13). Further confirmation of this trend was reported recently by Penprase et al. (2010), who proposed that the C/O ratio continued to rise to super-solar values at even lower metallicity based on measurements from 5 DLAs (see Figure 2.14). Their observations were, however, taken with the echelle spectrograph and imager (ESI) mounted on the Keck II telescope. Since ESI is a medium spectral resolution instrument ($R \approx 5000; \nu_{\text{FWHM}} \approx 60 \text{ km s}^{-1}$), these authors were restricted to a COG analysis, which has its limitations (see Section 2.3). In particular, line saturation and the multi-component structure exhibited by some metal-poor DLAs can be easily overlooked in medium resolution data.

Thus, although this result certainly seems plausible, it should first be confirmed with high spectral resolution data, where the DLA’s line profiles and component structure can be resolved, and the effects of saturation can be properly investigated. Certainly, these two concerns require the most attention when one measures the carbon abundance from the CII lines at 1036 Å and 1334 Å for several reasons. First, since the carbon abundance is measured from a singly ionised atom, there might exist some nearby, mildly ionised gas that gives rise to some additional CII...
absorption that is not associated with the neutral gas attributed to the DLA (traced very well by the O I lines). Second, both C II transitions give rise to absorption lines that are intrinsically very strong, and are therefore saturated in most DLAs.

These difficulties, in combination with the rarity of metal-poor DLAs, have therefore resulted in very few systems where one can accurately measure the carbon abundance in near pristine environments. However, measuring the carbon abundance is of utmost importance for a number of reasons, since carbon is (1) the first metal synthesised in the long chain of stellar nucleosynthesis; (2) thought to be the most abundant metal during the first episodes of metal enrichment; and (3) the dominant coolant of near-pristine gas, which later allowed the first low-mass stars to form. Therefore, measuring [C/H] is perhaps the most profitable step forward to build on our understanding about the nature of the first stars, and the transition between the epoch of massive Population III star formation to the one we observe today that is dominated by low-mass Population II stars. This thesis, entitled ‘Finding the First Metals’, could therefore be considered as a search for the first metal – carbon.
2.5 Organisational Notes

This thesis is comprised of 7 Chapters, which mostly focus on the goal of searching for the first metals that were synthesised in our Universe. Chapter 3 reports the typical sizes and star formation rates of DLAs, whilst Chapters 4, 5, and 6 focus on using the most metal-poor DLAs as probes of early nucleosynthesis. In addition, Chapter 6 outlines how this work has contributed to our understanding of the first stars, and the future work that can be built on the findings I report herein. Finally, Chapter 7 reports additional, unrelated work that was also completed during my Ph.D. Chapters 3, 5, and 7 have all been accepted for publication in the *Monthly Notices of the Royal Astronomical Society*, whilst Chapter 4, is currently under review.

**Chapter 3: The sizes and star formation rates of DLAs**

I report the discovery of a DLA that is only observed along one line of sight to the gravitationally lensed quasar pair UM 673. In the neighbouring sightline I report a rare detection of Ly\(\alpha\) emission. When considered with additional examples in the literature, these discoveries are used to constrain the typical sizes and star formation rates of DLAs. This work is published as MNRAS 409, 679 (2010) with my coauthors M. Pettini, C. C. Steidel, L. J. King, G. C. Rudie, and O. Rakic.

**Chapter 4: A survey for the most metal-poor DLAs**

This Chapter presents the results from my survey to uncover the most metal-poor DLAs. This survey doubles the current number of known systems reported in the literature, and tentatively suggests that these systems might have received their metals from the explosions of Population III stars. The work in this Chapter is written in the format of a paper, and has been submitted to MNRAS for review. This work was completed with the assistance of M. Pettini, C. C. Steidel, G. C. Rudie, and P. E. Nissen.

**Chapter 5: Probing the nucleosynthesis from the first stars**

This Chapter reports a detailed study of the abundance pattern for a DLA towards one of the quasars presented in Chapter 4 whose abundance pattern cannot be explained by Population II nucleosynthesis, but is explained quite naturally with models of Population III nucleosynthesis. This work is published as MNRAS 412, 1047 (2011) with my coauthors M. Pettini, C. C. Steidel, G. C. Rudie, and R. A. Jorgenson.
Chapter 6: Finding the first metals

Finally, I conclude this topic by summarising the main results that have come from this work, and include discussion on the avenues now available for future research. Specifically, I discuss possible ways that will permit an easier identification of carbon-enhanced DLAs, and the role such systems will play in our understanding of the nature of the first stars. I speculate that some diagnostic ratios will provide information on the explosion energy, and degree of mixing in the metal-free stars.

Chapter 7: Does the Universe expand (an)isotropically?

During my Ph.D., I also had the opportunity to work on a recent compilation of distance estimates from Type Ia supernovae, which are believed to be excellent standard candles. In this Chapter, I test whether or not there is a dark energy dipole in our Universe. I conclude this Chapter by reporting a dipole that is directed towards that of the cosmic microwave background, although at very low significance. This work is published as MNRAS 401, 1409 (2010) with my co-author D. Lynden-Bell. Finally, I suggest some avenues for future research with type Ia supernovae that may allow us to study the presence or absence of a dark energy dipole in greater detail.
The sizes and star formation rates of DLAs

**Summary**

The sightline to the brighter member of the gravitationally lensed quasar pair UM 673A,B intersects a damped Lyα system at $z_{\text{abs}} = 1.62650$ which, because of its low redshift, has not been recognised before. The high quality echelle spectra of the pair, obtained with HIRES on the Keck I telescope, show a drop in neutral hydrogen column density $N(\text{H}I)$ by a factor of at least 400 between UM 673A and B, indicating that the DLA’s extent in this direction is much less than the $2.7h^{-1}_{70}$ kpc separation between the two sightlines at $z_{\text{abs}} = 1.62650$. By reassessing this new case together with published data on other quasar pairs, I conclude that the typical size (radius) of DLAs at these redshifts is $R \approx (5 \pm 3)h^{-1}_{70}$ kpc, smaller than previously realised. Highly ionized gas associated with the DLA is more extended, as there are only small differences in the C IV absorption profiles between the two sightlines.

Coincident with UM 673B, I detect a weak and narrow Lyα emission line which I attribute to star formation activity at a rate $\text{SFR} \gtrsim 0.2 \text{M}_\odot \text{yr}^{-1}$. The DLA in UM 673A is metal-poor, with an overall metallicity $Z_{\text{DLA}} \approx 1/30Z_\odot$, and has a very low internal velocity dispersion. It exhibits some apparent peculiarities in its detailed chemical composition, with the elements Ti, Ni, and Zn being deficient relative to Fe by factors of 2–3. The [Zn/Fe] ratio is lower than those measured in any other DLA or Galactic halo star, presumably reflecting somewhat unusual previous enrichment by stellar nucleosynthesis. I discuss the implications of these results for the nature of the galaxy hosting the DLA.
3.1 Introduction

UM 673 (Q0142−100) was first identified as a gravitationally lensed quasar at $z_{\text{em}} = 2.719$ by Surdej et al. (1987) who showed it to be lensed by a $z = 0.49$ galaxy into two images, UM 673A and UM 673B, separated by 2.2 arcsec and with magnitudes $m_R = 16.9$ and 19.1 respectively (see Figure 3.1). This pair was later examined in detail by Smette et al. (1992) in their study of intergalactic Ly$\alpha$ forest absorbers. By considering the number of absorption lines in common (and not in common) between the two closely spaced sightlines, these authors were able to place interesting limits on the typical sizes of Ly$\alpha$ clouds, following on from similar analyses by Sargent et al. (1982) and Foltz et al. (1984).

![Figure 3.1: Left panel: An image of UM 673A (right quasar) and UM 673B (left quasar). Right panel: After removing the lensed quasar, the foreground lensing galaxy comes into view.](image)

Since these early papers, there have been many investigations of the Ly$\alpha$ forest in the spectra of gravitationally lensed QSOs. However, very little is still known about the size and geometry of the neutral gas reservoirs that give rise to DLAs, since these are much rarer absorption systems and thus unlikely to be found in front of gravitationally lensed QSOs which are themselves unusual alignments. A recent study by Monier et al. (2009) of the quadruply lensed Cloverleaf QSO (H 1413+117) with the Hubble Space Telescope uncovered three new DLAs or sub-DLAs (defined to have $19.0 \leq \log[N(\text{H}I)/\text{cm}^{-2}] \leq 20.3$; Péroux et al. 2003a) at $z \sim 1.5$, none of which are common to all four components. When considered together with analogous observations of four other DLAs from the literature, these data led Monier et al. (2009) to conclude that absorbers at $z \sim 1.5$ with $N(\text{H}I) = (6 - 13) \times 10^{20} \text{cm}^{-2}$ have typical scale-lengths $S_{\text{DLA}} = (6 - 12) h_{70}^{-1} \text{kpc}$ (where $h_{70}$ is the Hubble constant in units of $70 \text{km s}^{-1} \text{Mpc}^{-1}$). Apparently at odds with this conclusion is the finding by Ellison et al. (2007) of coincident damped Lyman-α absorption on 100 kpc-scale towards the binary QSO SDSS 1116+4118A,B. As these dimensions far exceed those expected for a single galaxy, Ellison et al. (2007) favour an explanation in terms of a group of two or more galaxies intersected along these lines of sight.

In this Chapter I report the discovery of a previously unrecognised DLA, at $z_{\text{abs}} = 1.62650$, in the spectrum of UM 673A from high resolution and high signal-to-noise ratio (S/N) spectroscopy with the HIRES instrument on the Keck I telescope. In line with the compilation of
similar measurements by Monier et al. (2009), no DLA is seen in front of UM 673B, even though at the redshift of the DLA the two sightlines are separated by less than 3 kpc.

There is, however, a weak Lyα emission line in the spectrum of UM 673B at the same redshift as the DLA. There have been only a few reported cases of Lyα emission associated with a DLA since the first such detection by Hunstead et al. (1990), as summarised in the review of DLA properties by Wolfe et al. (2005) with more recent updates by Kulkarni et al. (2006) and Christensen et al. (2007). Any new examples are of interest in view of the apparent puzzle presented by the lack of obvious star formation associated with gas-rich DLAs (Wolfe & Chen, 2006), and the recent claim that a newly discovered population of faint line emitters represents the long-sought host galaxies of DLAs (Rauch et al., 2008).

Finally, the newly discovered DLA in UM 673A is metal-poor, with metallicity $Z \sim 1/30Z_\odot$. Such chemically unevolved DLAs are important, in that they can provide clues to early episodes of metal enrichment in the Universe, complementing efforts being directed to analogous studies of metal-poor stars in the Milky Way and nearby dwarf galaxies (see, for example, Pettini, 2006). In the present case, I measure the relative abundances of nine different elements, from N to Zn, and uncover some chemical peculiarities which have not been noted before.

This Chapter is organised as follows. In Section 3.2, I briefly describe the observations of UM 673A,B and the reduction of the HIRES spectra. In the subsequent analysis, I first focus on the H\textsc{i} gas in front of the UM 673 pair (Section 3.3), and consider the data presented here together with the compilation by Monier et al. (2009) of other DLAs in gravitationally lensed QSOs to refine those authors' estimate of the characteristic size of damped systems (Section 3.4). I next turn to the Lyα emission detected in the spectrum of UM 673B at the same redshift as the DLA in UM 673A (Section 3.5), and use it to obtain an estimate of the star formation rate in the galaxy associated with the DLA. Section 3.6 deals with the chemical composition of the DLA, comparing it to that of DLAs and Galactic halo stars of similar overall metallicity. I summarise my findings and draw some conclusions in Section 3.7. Throughout this Chapter, I adopt a ‘737’ cosmology, with $H_0 = 70\text{ km s}^{-1}\text{ Mpc}^{-1}$, $\Omega_M = 0.3$ and $\Omega_\Lambda = 0.7$.

### 3.2 Observations and Data Reduction

UM 673A and UM 673B were observed on the nights of 2005 October 9 and 10, and again three years later on the nights of 2008 September 24 and 25, as part of a large-scale imaging and spectroscopic survey of galaxies in the fields of bright QSOs, aimed primarily at studying the outflows of interstellar gas from star-forming galaxies at redshifts $z = 2–3$ (Adelberger et al., 2005; Steidel et al., 2010). HIRES (Vogt et al., 1994) was configured to cover the wavelength range 3100–6100 Å (with small gaps near 4000 Å and 5000 Å due to gaps between the three CCD chips on the detector) using the ultraviolet (UV) cross-disperser and collimator.

In order to avoid cross-contamination between the two images and to minimise slit losses
due to atmospheric dispersion, UM 673 A and B were observed separately, with the HIRES slit maintained at the parallactic angle by its image rotator. For UM 673 A a 1.15 arcsec-wide entrance slit was employed, which results in a resolution $R \equiv \lambda / \Delta \lambda = 36000$, corresponding to a velocity full width at half maximum FWHM = 8.3 km s$^{-1}$, sampled with $\sim$ 3 pixels. The total integration time was 9400 s, divided into five exposures; the QSO was stepped along the slit between each exposure. For the fainter UM 673 B, a narrower 0.86 arcsec slit was used (so as to exclude more effectively light from the brighter image) which results in $R = 48000$ and FWHM = 6.2 km s$^{-1}$ sampled with $\sim$ 2 detector pixels. The total exposure time was 28200 s, again divided into a number of separate exposures, typically 2700 s long. The seeing was $\leq$ 1 arcsec FWHM throughout the observations.

To these data I added another set of observations of UM 673A,B, which I retrieved from the Keck Observatory data archive, obtained in 1996 with the original HIRES red-sensitive detector and red-optimised cross-disperser. While these earlier data do not contribute much at blue and ultraviolet wavelengths, with their long exposure times (18 000 s and 27 000 s for UM 673A and B respectively) they do improve the S/N ratio of the final co-added spectrum at red wavelengths.

The two-dimensional HIRES spectra were processed with the *MAKEE* data reduction pipeline developed by Tom Barlow which includes the usual steps of flat-fielding, order tracing, background subtraction, 1-D extraction and merging of the echelle orders. A wavelength reference was provided by the spectrum of the internal Th-Ar hollow cathode lamp and the co-added, 1-D spectra were mapped onto a vacuum heliocentric wavelength scale. In a final step, the spectra were normalised by dividing out the QSO continuum and emission lines. The rms deviations of the data from the continuum in regions free from absorption lines provide an empirical measure of the signal-to-noise ratio. For UM 673A, the data have S/N $> 35$ per 2.7 km s$^{-1}$ ($\approx$ 0.04 Å) bin from $\sim$ 4000 Å to $\sim$ 6000 Å; the S/N is highest near 5000 Å (S/N $\approx$ 70) and is still moderately high (S/N $\approx$ 24) at 3200 Å, near the redshifted wavelength of the damped Ly$\alpha$ line. The corresponding values for UM 673B are S/N $\gtrsim$ 15 (4000–6000 Å) and S/N $\approx$ 8 at 3200 Å.

### 3.3 HI Absorption towards UM 673A,B

#### 3.3.1 The DLA towards UM 673A

These HIRES spectra extend to shorter wavelengths than most previous observations of this famous QSO pair, which probably explains why the damped Ly$\alpha$ system in front of UM 673A (see Figure 3.2) has gone unnoticed until now. Associated with the DLA are a multitude of metal absorption lines of elements from C to Zn. These lines are analysed in detail in Section 3.6; for the present purpose suffice it to say that they are narrow, with FWHM $\lesssim 25$ km s$^{-1}$, and have maximum optical depth at $z_{\text{abs}} = 1.626498$. Adopting the same redshift for the damped Ly$\alpha$ line, I find $\log [N(\text{H}I)/\text{cm}^{-2}] = 20.7 \pm 0.1$ by fitting theoretical Voigt profiles to the wings of the line (and interpolating across narrower absorption features — see Figure 3.2).
3.3 HI Absorption towards UM 673A,B

Figure 3.2: Portion of the HIRES spectrum of UM 673A (black histogram) encompassing the damped Lyα line at $z_{\text{abs}} = 1.626498$. The red continuous line shows the theoretical Voigt profile for a neutral hydrogen column density $\log [N(\text{H} I)/\text{cm}^{-2}] = 20.7$. The normalised continuum and zero-level are shown by the blue dashed and green dotted lines respectively. The $y$-axis scale is residual intensity.

3.3.2 The Lyman limit system in UM 673B

I also find absorption near $z_{\text{abs}} = 1.626498$ in the spectrum of UM 673B, but with much reduced column densities of neutral gas. The top panel of Figure 3.3 shows the wavelength region around the Lyα line which is broad and saturated, but not damped. Under these circumstances, it is well known that the column density is unconstrained within orders of magnitude, unless higher order Lyman lines are available—this is the reason why the H I column density distribution is so poorly sampled in the interval $\log [N(\text{H} I)/\text{cm}^{-2}] = 17–20$ (e.g. Storrie-Lombardi & Wolfe, 2000; O’Meara et al., 2007).

However, since $N(\text{H} I)$ and the velocity dispersion parameter $b$ (km s$^{-1}$) are degenerate in strongly saturated lines, I can still derive an upper limit to the column density by considering the smallest $b$-value, and corresponding highest value of $N(\text{H} I)$, which provide an acceptable fit to the width and profile of the saturated Lyα line in UM 673B. To this end, I considered a series of pair values of $b$ and $N(\text{H} I)$, fixing $N(\text{H} I)$ and using VPFIT\(^1\) to determine the value of $b$ for which the theoretical line profile shows the least disagreement with the data.

I started the iteration at $\log [N(\text{H} I)/\text{cm}^{-2}] = 20.7$, as measured in UM 673A and which greatly overproduces the observed Lyα absorption in UM 673B for all values of $b$, and then decreased $\log N(\text{H} I)$ in steps of 0.1, until a plausible fit was arrived at for $\log [N(\text{H} I)/\text{cm}^{-2}] = 18.1$ and $b = 22$ km s$^{-1}$. The corresponding line profile, convolved with the instrumental resolution, is superimposed on the data in the top panel of Figure 3.3. While the absorption in the line core is less than observed, presumably because of neighbouring Lyα absorption lines—one of which, at $\Delta v = +114$ km s$^{-1}$, is incidentally also seen as a redshifted component in C IV—higher column densities of $N(\text{H} I)$ would overproduce the absorption in the line wings, relative to what is observed. I consider $\log [N(\text{H} I)/\text{cm}^{-2}] \leq 18.1$ to be an upper limit to the column density of neutral gas in UM 673B because equally good or even better fits could be obtained.

\(^1\)VPFIT is available from http://www.ast.cam.ac.uk/~rfc/vpfit.html
Figure 3.3: A selection of absorption lines in UM 673B near the redshift of the DLA at $z_{\text{abs}} = 1.626498$ in front of UM 673A; in all panels the $y$-axis is residual intensity. The three green arrows in the top panel indicate the velocities of the three components of the low-ionisation metal lines in the DLA (see Section 3.6), while the three long-dash red lines through all the panels mark the velocities of absorption (and emission) components in UM 673B. In the top panel, the red continuous line is the theoretical Ly$\alpha$ absorption profile for the upper limit I deduce to the column density of H$\text{i}$ in UM 673B, $\log[N(\text{H}\text{i})/\text{cm}^{-2}] = 18.1$. The shaded area shows the $\pm1\sigma$ error spectrum. Note the detection of Ly$\alpha$ emission in the core of the strong absorption line, at the same redshift as the DLA in UM 673A. The emission profile also aligns well with the high ionisation absorption lines in UM 673B reproduced in the lower four panels.
3.4 Constraining the sizes of DLAs

with lower values of \( N(\text{H}^1) \) and higher values of \( b \).

Thus, I am led to the conclusion that the column density of neutral hydrogen drops by a factor of at least 400 over a transverse distance of less than 3 kpc (Section 3.4). A comparable drop is deduced from consideration of metal absorption lines from ionisation stages which are dominant in \( \text{H}^1 \) regions. For example, in Section 3.6, I deduce a column density \( \log[N(\text{Si}^\text{II})/\text{cm}^{-2}] = 14.75 \pm 0.03 \) from the analysis of five \( \text{Si}^\text{II} \) transitions in UM 673A. The strongest of these, \( \text{Si}^\text{II} \lambda 1260.4221 \), is below the detection limit in UM 673B (see second panel from the top in Figure 3.3). In the optically thin limit,

\[
N = 1.13 \times 10^{20} \cdot \frac{W_\lambda}{\lambda^2 f} \text{ cm}^{-2} \tag{3.1}
\]

where \( W_\lambda \) and \( \lambda \) are respectively the rest frame equivalent width and wavelength (both in Å), and \( f \) is the oscillator strength.\(^2\) From equation (3.1) I deduce \( \log[N(\text{Si}^\text{II})/\text{cm}^{-2}] \leq 12.2 \) (3σ limit) in UM 673B, a factor of \( \geq 350 \) lower than in UM 673A.

### 3.4 Constraining the sizes of DLAs

In this section I use the finding that the DLA in UM 673A is not present in the spectrum of UM 673B to reassess, together with existing data, the characteristic size of the \( \text{H}^1 \) clouds giving rise to damped \( \text{Ly}^\alpha \) systems.

I begin with a simple derivation of the transverse distance, similar to that presented by Smette et al. (1992). Referring to Figure 3.4, the transverse distance between the two images at the lens plane is \( S_L = \theta_{\text{obs}} D_{\text{OL}} = \alpha D_{\text{SL}}, \) and the transverse distance between the two light paths at the redshift of the absorber is \( S_0 = \alpha D_{\text{SC}}, \) where \( D_{\text{OL}}, D_{\text{SL}}, D_{\text{SC}} \) are the angular diameter distances from, respectively, the observer to the lens, the source to the lens, and the source to the absorbing cloud.

Thus, the transverse distance between the light paths at the redshift of the absorber can be written as

\[
S_0 = \frac{\theta_{\text{obs}} D_{\text{OL}} D_{\text{SC}}}{D_{\text{SL}}} = \frac{\theta_{\text{obs}} D_{\text{OL}} D_{\text{CS}}}{D_{\text{LS}}} \cdot \frac{(1 + z_L)}{(1 + z_C)} \tag{3.2}
\]

where \( z_L \) and \( z_C \) are the redshifts of the lens and the DLA respectively. Recalling that the angular diameter distance between two objects at redshift \( z_2 \) and \( z_1, \) where \( z_2 > z_1, \) is of the form (Hogg, 1999),

\[
D_{12} = \frac{D_2 - D_1}{1 + z_2}, \tag{3.3}
\]

\(^2\)Throughout this work, I use the compilation of laboratory wavelengths and \( f \)-values by Morton (2003) with updates by Jenkins & Tripp (2006).
Figure 3.4: The geometry of a gravitational lens, as viewed by the observer at O, with source (S) being lensed into two images (I₁ and I₂) of angular separation $\theta_{\text{obs}}$, by a galaxy situated at the lens plane along the optical axis (dashed horizontal line). $S_0$ is the transverse distance between the two sightlines at the location of the DLA.

where

$$D_i = \frac{c}{H_0} \int_0^{z_i} \frac{dz}{\sqrt{\Omega_\Lambda + (1 + z)^3 \Omega_M}}$$

(3.4)

is the comoving distance to redshift $z_i$, I can rewrite equation (3.2) as:

$$S_0 = \frac{\theta_{\text{obs}} D_L (D_S - D_C)}{(1 + z_C)(D_S - D_L)}.$$

(3.5)

Thus, adopting $\theta_{\text{obs}} = 2.22$ arcsec, $z_L = 0.493$ (Surdej et al., 1988; Lehár et al., 2000), $z_S = 2.7434$ from unpublished near-infrared observations of the Hβ emission line\(^3\), and $z_C = 1.62650$ from the HIRES observations presented here, I find that the transverse (physical) distance between the two sightlines at the redshift of the DLA is $S_0 = 2.7$ kpc (in the adopted ‘737’ cosmology).

I now add this new case to the list compiled by Monier et al. (2009) and reanalyse the entire sample in Table 3.1. Note that I have revised the values for the Cloverleaf (H 1413+117) from Table 5 of Monier et al. (2009) for two reasons. First, in calculating the transverse distances applicable to the two low redshift systems (at $z_{\text{abs}} = 1.440$ and 1.486) in the Cloverleaf, these authors assumed $z_L = z_C$ (E. Monier, private communication). Second, recent mid-infrared data have improved the understanding of this system (MacLeod et al., 2009). The positions and relative fluxes of the images can be well explained by a lensing galaxy at $z_L \approx 1.0$ (Kneib, Alloin, & Pello, 1998), with an additional galaxy located $\Delta \alpha_{G2} = -1.87$ arcsec, $\Delta \delta_{G2} = 4.14$ arcsec from the lens, coincident with a galaxy identified by Kneib et al. (1998) as object No. 14. The inclusion of this additional galaxy in the lensing model does not affect the astrometry (but does affect the relative fluxes of the QSO images); thus, I assume a single lens geometry at $z = 1.0$. I

\(^3\)As is normally the case, the systemic redshift deduced from the Balmer lines is higher than the redshift indicated by the rest-frame ultraviolet emission lines, in this case $z_{\text{em}} = 2.719$ from the EFOSC spectra obtained by Surdej et al. (1987) and $z_{\text{em}} = 2.7313$ from the Sloan Digital Sky Survey (Schneider et al., 2007). The difference from the original value reported by Surdej et al. (1987) is nearly +2000 km s\(^{-1}\).
### Table 3.1: Derived e-folding scale lengths of DLAs, updated from Table 5 of Monier et al. (2009)

<table>
<thead>
<tr>
<th>QSO</th>
<th>$z_{em}$</th>
<th>$z_{lens}$</th>
<th>$z_{abs}$</th>
<th>Pair</th>
<th>$\theta_{obs}$ (arcsec)</th>
<th>$N(\text{H}I)_{\text{max}}$ (10^{20} \text{ cm}^{-2})$</th>
<th>$N(\text{H}I)_{\text{min}}$ (10^{20} \text{ cm}^{-2})$</th>
<th>$S^e_0$ ($h^{-1}_{70}$ \text{kpc})</th>
<th>$S^e_{\text{DLA},e}$ ($h^{-1}_{70}$ \text{kpc})</th>
<th>$S^e_{\text{DLA},e}$ ($h^{-1}_{70}$ \text{kpc})</th>
</tr>
</thead>
<tbody>
<tr>
<td>H1413+117</td>
<td>2.55</td>
<td>1.0</td>
<td>1.440</td>
<td>B-A</td>
<td>0.753</td>
<td>60</td>
<td>9.0</td>
<td>3.13</td>
<td>2.33</td>
<td>1.65</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>B-D</td>
<td>0.967</td>
<td>60</td>
<td>0.25</td>
<td>4.02</td>
<td>2.55</td>
<td>0.733</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>B-C</td>
<td>1.359</td>
<td>60</td>
<td>0.20</td>
<td>5.65</td>
<td>3.58</td>
<td>0.991</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>A-D</td>
<td>1.118</td>
<td>9.0</td>
<td>0.25</td>
<td>4.64</td>
<td>3.02</td>
<td>1.29</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>A-C</td>
<td>0.872</td>
<td>9.0</td>
<td>0.20</td>
<td>3.62</td>
<td>2.34</td>
<td>0.951</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>D-C</td>
<td>0.893</td>
<td>0.25</td>
<td>0.20</td>
<td>3.71</td>
<td>11.7</td>
<td>16.6</td>
</tr>
<tr>
<td>H1413+117</td>
<td>1.486</td>
<td></td>
<td></td>
<td>D-A</td>
<td>1.118</td>
<td>2.0</td>
<td>&lt;0.05</td>
<td>4.32</td>
<td>&lt;2.80</td>
<td>&lt;1.17</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>D-B</td>
<td>0.967</td>
<td>2.0</td>
<td>&lt;0.1</td>
<td>3.73</td>
<td>&lt;2.48</td>
<td>&lt;1.25</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>D-C</td>
<td>0.893</td>
<td>2.0</td>
<td>&lt;0.05</td>
<td>3.45</td>
<td>&lt;2.24</td>
<td>&lt;0.935</td>
</tr>
<tr>
<td>H1413+117</td>
<td>1.662</td>
<td></td>
<td></td>
<td>B-A</td>
<td>0.753</td>
<td>6.0</td>
<td>1.5</td>
<td>2.17</td>
<td>1.83</td>
<td>1.57</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>B-C</td>
<td>1.359</td>
<td>6.0</td>
<td>0.6</td>
<td>3.91</td>
<td>2.75</td>
<td>1.70</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>B-D</td>
<td>0.967</td>
<td>6.0</td>
<td>0.3</td>
<td>2.78</td>
<td>1.85</td>
<td>0.928</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>A-C</td>
<td>0.872</td>
<td>1.5</td>
<td>0.6</td>
<td>2.51</td>
<td>2.64</td>
<td>2.74</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>A-D</td>
<td>1.118</td>
<td>1.5</td>
<td>0.3</td>
<td>3.22</td>
<td>2.54</td>
<td>2.00</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td></td>
<td>C-D</td>
<td>0.893</td>
<td>0.6</td>
<td>0.3</td>
<td>2.57</td>
<td>3.25</td>
<td>3.71</td>
</tr>
<tr>
<td>HE0512−3329</td>
<td>1.58</td>
<td>0.93</td>
<td>0.9313</td>
<td>A-B</td>
<td>0.644</td>
<td>3.09</td>
<td>2.95</td>
<td>5.05</td>
<td>70.4</td>
<td>109</td>
</tr>
<tr>
<td>Q0957+561</td>
<td>1.4136</td>
<td>0.36</td>
<td>1.3911</td>
<td>A-B</td>
<td>6.2</td>
<td>1.9</td>
<td>0.8</td>
<td>0.278</td>
<td>0.303</td>
<td>0.321</td>
</tr>
<tr>
<td>UM673</td>
<td>2.7313</td>
<td>0.493</td>
<td>1.6265</td>
<td>A-B</td>
<td>2.22</td>
<td>5.0</td>
<td>&lt;0.013</td>
<td>2.71</td>
<td>&lt;1.72</td>
<td>&lt;0.455</td>
</tr>
<tr>
<td>HE1104−1805</td>
<td>2.31</td>
<td>0.73</td>
<td>1.6616</td>
<td>A-B</td>
<td>3.0</td>
<td>6.3</td>
<td>&lt;0.037</td>
<td>4.47</td>
<td>&lt;2.84</td>
<td>&lt;0.870</td>
</tr>
</tbody>
</table>

*a* Transverse separation between the two sightlines at the redshift of the absorber.

*b* e-folding scale length of DLA assuming a linear decline of $N(\text{H}I)$—see equation (3.7).

*c* e-folding scale length of DLA assuming an exponential decline of $N(\text{H}I)$—see equation (3.9).
have excluded from the Monier et al. (2009) sample the binary QSO LBQS 1429−0053 because the linear scale probed by that pair, $S_0 = 43.2 h_{70}^{-1}$ kpc, is one order of magnitude larger than those of all the other (gravitationally lensed) pairs considered here. As argued by Ellison et al. (2007), such large scales are more likely to probe the clustering properties of DLAs rather than the typical sizes of their host galaxies. In any case, my conclusions below on the median size of DLAs are unaltered by the inclusion or omission of LBQS 1429−0053.

In order to determine the characteristic size of DLAs, $S_{DLA}$, I adopt two simple models. In the first, $S_{DLA}$ depends linearly on the $\text{HI}$ column density, while in the second $S_{DLA}$ depends linearly on the logarithm of $N(\text{HI})$. The latter is probably more realistic, as it corresponds to an exponential decline of $N(\text{HI})$, but I also consider the former for comparison with the analysis by Monier et al. (2009). My analysis, however, differs from theirs in the following way. Monier et al. (2009) examined $N(\text{HI})$ as a function of the observed angular separation of two QSO images, $\theta_{\text{obs}}$, whereas I consider $N(\text{HI})$ as a function of the physical (transverse) distance between the two sightlines at the redshift of the absorber. The approach taken by Monier et al. (2009) has the advantage of being independent of the choice of cosmological parameters, while mine is perhaps more physically motivated. We therefore have an equation of the form,

$$S = S_0 \frac{N(\text{HI})_{\text{max}} - N(\text{HI})}{N(\text{HI})_{\text{max}} - N(\text{HI})_{\text{min}}},$$

(3.6)

where $N(\text{HI})_{\text{max}}$ is the higher column density observed between any two given sightlines, and $N(\text{HI})_{\text{min}}$ is the lower column density of the two. In the first model, I follow Monier et al. (2009) and define the linear $e$-folding scale-length ($S_{DLA,\bar{e}}$) to be the transverse distance over which $N(\text{HI})_{\text{max}}$ decreases by a factor of $e = 2.718$ (i.e. $N(\text{HI}) = N(\text{HI})_{\text{max}}/2.718$),

$$S_{DLA,\bar{e}} = S_0 \frac{0.632 N(\text{HI})_{\text{max}}}{N(\text{HI})_{\text{max}} - N(\text{HI})_{\text{min}}}.$$

(3.7)

Note that one can easily convert to the ‘$e$-folding angle’, $\theta_{\text{e}}$ introduced by Monier et al. (2009) using the relation: $\theta_{\text{e}} = \theta_{\text{obs}} S_{DLA,\bar{e}}/S_0$.

In the second case considered, $S$ scales with $\ln N(\text{HI})$:

$$N(\text{HI}) = N(\text{HI})_{\text{max}} \exp[-S/S_{DLA,e}].$$

(3.8)

I then define the true $e$-folding scale length for DLAs,

$$S_{DLA,e} = \frac{S_0}{\ln[N(\text{HI})_{\text{max}}/N(\text{HI})_{\text{min}}]},$$

(3.9)

where $S_0$, $N(\text{HI})_{\text{max}}$ and $N(\text{HI})_{\text{min}}$ all take their previous definitions.

In both models, the $e$-folding scale lengths are most uncertain for $N(\text{HI})_{\text{max}}/N(\text{HI})_{\text{min}} \approx 1$. 
3.4 Constraining the sizes of DLAs

Figure 3.5: The distributions of (a) linear $e$-folding scale lengths and (b) $e$-folding scale lengths for DLAs, based on equations (3.7) and (3.9) respectively. Upper limits in the values of $\tilde{S}_{DLA,\bar{e}}$ and $\tilde{S}_{DLA,e}$ in Table 3.1 have been plotted as if they were measured values. For presentation purposes, two outlying values in excess of 10 kpc have been omitted from these plots.

due to the large extrapolation required. The derived $e$-folding scale lengths for each DLA are listed in Table 3.1, and the two distributions are shown with histograms in Figure 3.5. Considering all the measurements, I determine the median linear $e$-folding and median $e$-folding scale lengths of DLAs to be $\tilde{S}_{DLA,\bar{e}} = 2.6 \pm 0.7$ kpc and $\tilde{S}_{DLA,e} = 1.3 \pm 0.8$ kpc respectively. The errors were computed with the Interactive Data Language routine ROBUST_SIGMA\(^4\) which determines the median absolute deviation (unaffected by outliers) of a set of measurements, and then appropriately weights the data to provide a robust estimate of the sample dispersion (Hoaglin, Mosteller, & Tukey, 1983).

The values of $\tilde{S}_{DLA,\bar{e}}$ and $\tilde{S}_{DLA,e}$ I deduce are lower than that reported by Monier et al. (2009), $S_{DLA} = 6 h^{-1}_{70}$ kpc, partly because of the improved estimate of the lens redshift in the Cloverleaf. Indeed, given the large number of sightline pairs in this multiple system (see Table 3.1), the uncertainty in the lens redshift of the Cloverleaf still has a marked effect on the values of $\tilde{S}_{DLA,\bar{e}}$ and $\tilde{S}_{DLA,e}$ deduced. In order to assess the effect quantitatively, I have repeated the above analysis with the rather extreme assumptions that the lens redshift is, in turn, $z_L = 0.5$ and $1.5$, instead of the value $z_L = 1.0$ adopted in Table 3.1 from Kneib et al. (1998). I find $\tilde{S}_{DLA,\bar{e}} = 1.0 \pm 0.3$ kpc, $\tilde{S}_{DLA,e} = 0.5 \pm 0.3$ kpc if $z_L = 0.5$, and $\tilde{S}_{DLA,\bar{e}} = 5.6 \pm 1.9$ kpc, $\tilde{S}_{DLA,e} = 2.7 \pm 2.0$ kpc if $z_L = 1.5$. The range of values of $S_{DLA}$ admitted by the data will narrow as more QSO pairs are studied in the future.

Finally, I stress that my estimates of $S_{DLA}$ are not the same as what is generally thought of as the ‘size’ of a DLA. Referring to equation (3.8), if I assume an idealised spherical cloud with peak $N(HI)_{\text{max}} = 1 \times 10^{22}$ cm$^{-2}$ at its centre, we have to move a radial distance $S = \ln 50 S_{DLA,e}$, or $S \simeq 4 S_{DLA,e}$, before $N(HI)$ falls below the threshold $N(HI)_{\text{min}} = 2 \times 10^{20}$ cm$^{-2}$ generally adopted as the definition of a DLA. Thus, my preferred solution, $\tilde{S}_{DLA,e} = 1.3 \pm 0.8$ kpc, corresponds to DLA radii of $\sim 5 \pm 3$ kpc.

\(^4\)Available from http://idlastro.gsfc.nasa.gov/homepage.html
Chapter 3. The sizes and star formation rates of DLAs

Figure 3.6: SDSS spectrum of UM 673A,B (black line) together with the power-law fit to the continuum (red line) of the form \( F_\lambda = A \cdot (\lambda / \AA)^{-\beta} \), with best fitting values \( \beta = 1.535 \) and \( A = 4.0 \times 10^{-10} \text{erg s}^{-1} \text{cm}^{-2} \text{Å}^{-1} \).

3.5 Ly\(\alpha\) emission towards UM 673B

Returning to Figure 3.3 (top panel), the presence of a weak emission line in the core of the saturated Ly\(\alpha\) absorption line can be readily recognised. Although in principle this feature could also be a gap between two adjacent absorption lines, its precise wavelength match with: (a) high ionisation metal absorption lines along the same sightline (lower four panels in Figure 3.3), and (b) the highest optical depth absorption in the DLA in UM 673A argues in favour of the interpretation as a weak and narrow Ly\(\alpha\) emission line.

In order to measure the line flux and luminosity, I referred the echelle spectrum (for which the absolute flux calibration can be uncertain) to the Sloan Digital Sky Survey (SDSS) spectrum of UM 673A,B, reproduced in Figure 3.6. Fitting the continuum longwards of the QSO Ly\(\alpha\) emission line with a power-law of the form \( F_\lambda = A \cdot (\lambda / \AA)^{-\beta} \), I deduced the best-fitting values \( \beta = 1.535 \pm 0.005 \) and \( A = (4.0 \pm 0.2) \times 10^{-10} \text{erg s}^{-1} \text{cm}^{-2} \text{Å}^{-1} \) for the slope and normalisation respectively. The fit, which is shown with a red line in Figure 3.6, reproduces the sum of the SDSS \( r \) magnitudes of UM 673A,B (\( r = 16.73 \) and 18.84, respectively) when convolved with the transmission curve of the \( r \)-band filter. Extrapolating this continuum to \( \lambda_{\text{obs}} = 3193 \text{Å} \), where the redshifted Ly\(\alpha\) line at \( z = 1.62650 \) falls, and allowing for the fraction of the light contributed by UM 673A, then provides an absolute flux scale for the continuum shown by the long-dash line in the top panel of Figure 3.3, where a residual intensity of 1.0 corresponds to a flux density \( F_B(3193\text{Å}) = (2.1 \pm 0.4) \times 10^{-16} \text{erg s}^{-1} \text{cm}^{-2} \text{Å}^{-1} \). The 20% error is the systematic uncertainty in the flux calibration due to the combined effects of: (i) extrapolation of the QSO continuum to 3193 Å, and (ii) the accuracy of the SDSS photometry.

Integrating over the Ly\(\alpha\) emission line then yields a line flux \( F(\text{Ly}\alpha) = (2.5 \pm 0.25 \pm 0.5) \times 10^{-17} \text{erg s}^{-1} \text{cm}^{-2} \), quoting separately the random error from the counting statistics (shaded region in the top panel of Figure 3.3) and the systematic uncertainty in the flux calibration. The Ly\(\alpha\) line luminosity in the adopted cosmology is \( L(\text{Ly}\alpha) = (4.3 \pm 0.4 \pm 0.9) \times 10^{41} \text{erg s}^{-1} \).
3.5 Lyα emission towards UM 673B

3.5.1 Origin of the Lyα emission

I now discuss some of the mechanisms that could produce the observed Lyα emission, before detailing what I consider to be the most plausible interpretation. The first source I consider is the metagalactic UV background, which could produce Lyα emission by fluorescence, the so-called Hogan-Weymann effect (Hogan & Weymann, 1987). This diffuse emission has yet to be observed, presumably because of the low surface brightness it produces. Given the current understanding of this effect, at the redshift of the DLA I would expect to observe a surface brightness \( \mu \approx 5.7 \times 10^{-19} \left( \frac{\eta_E}{0.5} \right) \left( \frac{J}{5.1 \times 10^{-22}} \right) \text{ erg s}^{-1} \text{ cm}^{-2} \text{ arcsec}^{-2}, \) (3.10)

where \( \eta_E \) represents the efficiency with which the incident UV background is re-radiated as fluorescent Lyα emission and \( J \) is the ionising background at the Lyman limit at \( z \sim 2 \) (Bolton et al., 2005). Through the area of sky covered by the observations (0.86 × 3 arcsec, the latter being the size of the aperture used to extract the 1-D spectra from the raw 2-D HIRES images), the surface brightness of Eq. (3.10) would produce an integrated line flux \( F_{\text{Ly}\alpha} \sim 1.5 \times 10^{-18} \text{ erg s}^{-1} \text{ cm}^{-2}, \) one order of magnitude lower than the flux recorded.

Lyα fluorescence can also be induced by the UV radiation from a nearby AGN impinging on the DLA (e.g. Adelberger et al., 2006). Following the formalism introduced by Cantalupo et al. (2005), a source with monochromatic luminosity \( L_\nu(v) = L_\nu(v_{\text{LL}})(v/v_{\text{LL}})^{-\alpha} \), where \( h v_{\text{LL}} = 13.6 \text{ eV} \), at a physical distance \( r \) from the DLA will correspond to a “boost factor”

\[
B = 15.2 \frac{L_\nu(v_{\text{LL}})}{10^{30} \text{ erg s}^{-1} \text{ Hz}^{-1}} 0.7 \left( \frac{r}{1 \text{ Mpc}} \right)^{-2},
\] (3.11)

where, for self-shielded clouds, \( B \) is empirically related to the increase in the observed Lyα surface brightness relative to that induced by the metagalactic UV background, \( \text{SB(Ly}\alpha)/\mu = (0.74 + 0.50B^{0.89}) \). Assuming that the observed Lyα emission is uniform across the area of sky covered by the observations (0.86 × 3 arcsec), the corresponding surface brightness is \( \text{SB(Ly}\alpha) = 9.7 \times 10^{-18} \text{ erg s}^{-1} \text{ cm}^{-2} \text{ arcsec}^{-2}, \) therefore requiring a boost factor \( B \approx 50 \).

Consider now a typical QSO at \( z = 1.5 - 2 \), with spectral index \( \alpha = 1.76 \) and \( \lambda L_\lambda(\lambda) = 1.5 \times 10^{46} \text{ erg s}^{-1} \) at 1100 Å (Telfer et al., 2002). One can extrapolate this typical luminosity to the Lyman limit using the relation

\[
L_\nu(v_{\text{LL}}) = \frac{\lambda^2}{c} L_\lambda(\lambda) \left( \frac{\lambda}{\lambda_{\text{LL}}} \right)^{2-\alpha} \approx 4 \times 10^{30} \text{ erg s}^{-1} \text{ Hz}^{-1}.
\] (3.12)

Substituting \( B \), \( L_\nu(v_{\text{LL}}) \) and \( \alpha \) into Eq. (3.11) yields a physical distance \( r = 700 \text{kpc} \). Thus, if the Lyα emission that is seen in the spectrum of UM 673B were produced by a nearby source of UV photons, such a source would need to be located within \( r \sim 700 \text{kpc} \) from the DLA towards
UM 673, or \( \lesssim 1.4 \) arcmin in projection on the sky. As I have not identified any such source in the deep galaxy survey of this area of sky, and none have been reported by others, I consider it unlikely that the observed Ly\( \alpha \) emission is due to fluorescence.

Other possibilities have been put forward (e.g. Dijkstra, Haiman, & Spaans, 2006), but all appear less likely than the most straightforward explanation that the Ly\( \alpha \) emission is produced by recombination in H\( \text{II} \) regions ionised by early-type stars in a galaxy presumably associated with the DLA. Adopting the Kennicutt (1998) relationship between star formation rate (SFR) and H\( \alpha \) luminosity, and assuming the ratio Ly\( \alpha \)/H\( \alpha \) \( \simeq 8.7 \) appropriate for case B recombination, yields an equation of the form,

\[
\text{SFR} \left( \frac{M_{\odot}}{\text{yr}^{-1}} \right) = 9.1 \times 10^{-43} L(\text{Ly}\alpha) \times \frac{1}{1.8} \left( \text{erg s}^{-1} \right).
\]

(3.13)

where the correction factor of 1/1.8 accounts for the flattening of the stellar initial mass function for masses below 1\( M_{\odot} \) (Chabrier, 2003) compared to the single power law of the Salpeter IMF assumed by Kennicutt (1998).

Thus, the inferred line luminosity, \( L(\text{Ly}\alpha) = 4.3 \times 10^{41} \) erg s\(^{-1} \) implies a star formation rate SFR \( \simeq 0.2 \) M\( \odot \) yr\(^{-1} \). In reality this is likely to be a lower limit, given the ease with which Ly\( \alpha \) photons are destroyed through resonant scattering in a dusty medium. In addition, the HIRES slit may have captured only a fraction of the Ly\( \alpha \) emission, if it is spatially more extended than the B image of UM 673 (a possibility which I cannot readily assess with the spectroscopic observations). However, I note that such low levels of star formation are not unusual for DLA host galaxies at \( z \lesssim 1 \) (Péroux et al., 2011).

### 3.5.2 Lensed Ly\( \alpha \) emission?

I next turn to the issue of where the Ly\( \alpha \) emitting region is located and whether it too may be lensed by the foreground galaxy at \( z = 0.493 \). Since the angular diameter distances from the lens to UM 673 (1130 Mpc) and from the lens to the absorption system (1036 Mpc) differ by only \( \sim 10\% \), one would expect the lensing geometry to be largely unchanged. Thus, presumably the Ly\( \alpha \) photons are also lensed by the foreground galaxy into a sister image near to, but offset from, the A sightline. The fact that emission is not seen in the core of the damped Ly\( \alpha \) absorption line in UM 673A suggests that the HIRES slit was not well placed to capture the A counterpart of the Ly\( \alpha \) emission. A more rigorous approach involves detailed modelling of the mass distribution of the foreground lensing galaxy, which unfortunately is not well constrained (Lehár et al., 2000).

I also note that the observed Ly\( \alpha \) flux may be slightly magnified, or demagnified, by the foreground galaxy, implying that the intrinsic Ly\( \alpha \) luminosity is different from that observed. However, a firm determination of the magnification factor is made difficult by the uncertain mass distribution of the lensing galaxy. Apart from improved modelling of the lens, progress in
the interpretation of the Lyα emission would be greatly facilitated by follow-up near-infrared integral field spectroscopy aimed at detecting the redshifted Hα emission line at $z = 1.62650$. Such observations would map out the full extent of the emission region, without the limitations imposed by single-slit spectroscopy.

3.6 Chemical Composition of the DLA in front of UM 673A

As is normally the case, a multitude of metal absorption lines are associated with the DLA towards UM 673A. With the wide wavelength coverage, high resolution and S/N ratio of the HIRES data, I detect 37 atomic transitions from the elements C to Zn in a variety of ionisation stages, from neutrals to triply ionised species, as detailed in Table 3.2. Although not all of these lines are available for abundance analysis (some being blended or saturated), I nevertheless have access to a great deal of information on the detailed chemical composition of this DLA. In this section, I analyse these data which may throw light on the chemical enrichment history of the galaxy giving rise to the DLA and provide clues to stellar nucleosynthesis at low metallicities.

3.6.1 Profile Fitting

Figure 3.7 shows a selection of metal absorption lines associated with the DLA. The absorption is evidently confined to a narrow velocity range, with even the strongest lines only extending over FWHM $\simeq 25–30$ km s$^{-1}$. In order to deduce values for the column density $N$ (cm$^{-2}$) and velocity dispersion parameter $b$ (km s$^{-1}$), I employed the absorption line profile fitting software VPFIT, which uses a $\chi^2$ minimisation technique to fit multiple Voigt profiles simultaneously to several atomic transitions and returns the best fitting values of $N$ and $b$ together with the associated errors (see, for example, Rix et al., 2007). The theoretical line profiles generated by VPFIT are superimposed on the data in Figure 3.7. I now consider in turn gas of low, intermediate, and high ionisation.

Low ion transitions

The Si II, Cr II, Fe II and Ni II lines can all be fitted with a minimum of three absorption components, with the parameters listed in Table 3.3. The highest optical depth is measured at $z_{\text{abs}} = 1.626498$ (component number 3 in Table 3.3, or C3 for short), which I therefore use as the zero point of the relative velocity scale for the DLA. A second component (component number 1, or C1 for short) at $z_{\text{abs}} = 1.626348$, or $\Delta v = -17.1$ km s$^{-1}$, can be readily recognised in the profiles of the weaker lines (e.g. Fe II $\lambda 2249$) in Figure 3.7. What is not immediately obvious from the Figure is that a third component (C2), at $z_{\text{abs}} = 1.626454$ and separated by only $\Delta v = -5.0$ km s$^{-1}$ from C3, is required to reproduce the asymmetric profiles and widths of the stronger lines. Note the small $b$-values, of order 1 km s$^{-1}$, deduced for C1...
Table 3.2: **Metal lines detected at the redshift of the DLA in UM 673A**

<table>
<thead>
<tr>
<th>Ion</th>
<th>Wavelength(^a) (Å)</th>
<th>(f^a)</th>
<th>(W^b_0) (mÅ)</th>
<th>(\delta W^b_0) (mÅ)</th>
</tr>
</thead>
<tbody>
<tr>
<td>C(\text{II})</td>
<td>1334.5323</td>
<td>0.1278</td>
<td>140</td>
<td>2</td>
</tr>
<tr>
<td>C(\text{IV})</td>
<td>1548.2041</td>
<td>0.1899</td>
<td>...(^c) 0.559</td>
<td>14.8 0.9</td>
</tr>
<tr>
<td>C(\text{IV})</td>
<td>1550.7812</td>
<td>0.09475</td>
<td>73</td>
<td>2</td>
</tr>
<tr>
<td>N(\text{I})</td>
<td>1199.5496</td>
<td>0.132</td>
<td>71</td>
<td>2</td>
</tr>
<tr>
<td>N(\text{I})</td>
<td>1200.2233</td>
<td>0.0869</td>
<td>67</td>
<td>2</td>
</tr>
<tr>
<td>N(\text{I})</td>
<td>1200.7098</td>
<td>0.0432</td>
<td>55</td>
<td>2</td>
</tr>
<tr>
<td>O(\text{I})</td>
<td>1302.1685</td>
<td>0.048</td>
<td>129</td>
<td>2</td>
</tr>
<tr>
<td>Al(\text{II})</td>
<td>1670.7886</td>
<td>1.740</td>
<td>...(^c) 18.5</td>
<td>14.8 0.9</td>
</tr>
<tr>
<td>Al(\text{III})</td>
<td>1854.71829</td>
<td>0.559</td>
<td>14.8 0.9</td>
<td></td>
</tr>
<tr>
<td>Al(\text{III})</td>
<td>1862.79113</td>
<td>0.278</td>
<td>7.7 0.8</td>
<td></td>
</tr>
<tr>
<td>Si(\text{II})</td>
<td>1190.4158</td>
<td>0.292</td>
<td>89</td>
<td>2</td>
</tr>
<tr>
<td>Si(\text{II})</td>
<td>1193.2897</td>
<td>0.582</td>
<td>120</td>
<td>2</td>
</tr>
<tr>
<td>Si(\text{II})</td>
<td>1260.4221</td>
<td>1.18</td>
<td>142</td>
<td>2</td>
</tr>
<tr>
<td>Si(\text{II})</td>
<td>1304.3702</td>
<td>0.0863</td>
<td>83.1</td>
<td>0.9</td>
</tr>
<tr>
<td>Si(\text{II})</td>
<td>1526.7070</td>
<td>0.133</td>
<td>118</td>
<td>1</td>
</tr>
<tr>
<td>Si(\text{II})</td>
<td>1808.01288</td>
<td>0.00208</td>
<td>24.6</td>
<td>0.8</td>
</tr>
<tr>
<td>Si(\text{III})</td>
<td>1206.500</td>
<td>1.63</td>
<td>...(^c)</td>
<td></td>
</tr>
<tr>
<td>Si(\text{IV})</td>
<td>1393.76018</td>
<td>0.513</td>
<td>...(^c)</td>
<td></td>
</tr>
<tr>
<td>Si(\text{IV})</td>
<td>1402.77291</td>
<td>0.254</td>
<td>...(^c)</td>
<td></td>
</tr>
<tr>
<td>S(\text{II})</td>
<td>1250.578</td>
<td>0.00543</td>
<td>...(^c)</td>
<td></td>
</tr>
<tr>
<td>S(\text{II})</td>
<td>1253.805</td>
<td>0.0109</td>
<td>21</td>
<td>1</td>
</tr>
<tr>
<td>S(\text{II})</td>
<td>1259.518</td>
<td>0.0166</td>
<td>...(^c)</td>
<td></td>
</tr>
<tr>
<td>Cr(\text{II})</td>
<td>2056.25693</td>
<td>0.103</td>
<td>21.6</td>
<td>0.9</td>
</tr>
<tr>
<td>Cr(\text{II})</td>
<td>2062.23610</td>
<td>0.0759</td>
<td>14.0</td>
<td>0.8</td>
</tr>
<tr>
<td>Cr(\text{II})</td>
<td>2066.16403</td>
<td>0.0512</td>
<td>10.3</td>
<td>0.9</td>
</tr>
<tr>
<td>Fe(\text{II})</td>
<td>1260.533</td>
<td>0.0240</td>
<td>...(^c)</td>
<td></td>
</tr>
<tr>
<td>Fe(\text{II})</td>
<td>1608.4509</td>
<td>0.0577</td>
<td>94.3</td>
<td>0.8</td>
</tr>
<tr>
<td>Fe(\text{II})</td>
<td>1611.20034</td>
<td>0.00138</td>
<td>...(^c)</td>
<td></td>
</tr>
<tr>
<td>Fe(\text{II})</td>
<td>2249.8768</td>
<td>0.001821</td>
<td>24</td>
<td>1</td>
</tr>
<tr>
<td>Fe(\text{II})</td>
<td>2260.7805</td>
<td>0.00244</td>
<td>32</td>
<td>1</td>
</tr>
<tr>
<td>Ni(\text{II})</td>
<td>1317.217</td>
<td>0.057</td>
<td>...(^c)</td>
<td></td>
</tr>
<tr>
<td>Ni(\text{II})</td>
<td>1370.132</td>
<td>0.056</td>
<td>8.4</td>
<td>0.9</td>
</tr>
<tr>
<td>Ni(\text{II})</td>
<td>1454.842</td>
<td>0.0323</td>
<td>...(^c)</td>
<td></td>
</tr>
<tr>
<td>Ni(\text{II})</td>
<td>1502.148</td>
<td>0.0133</td>
<td>...(^c)</td>
<td></td>
</tr>
<tr>
<td>Ni(\text{II})</td>
<td>1741.5531</td>
<td>0.0427</td>
<td>10.3</td>
<td>0.5</td>
</tr>
<tr>
<td>Ni(\text{II})</td>
<td>1751.9157</td>
<td>0.0277</td>
<td>6.0</td>
<td>0.6</td>
</tr>
<tr>
<td>Zn(\text{II})</td>
<td>2026.13709</td>
<td>0.501</td>
<td>4.8</td>
<td>0.7</td>
</tr>
</tbody>
</table>

\(^a\) Laboratory wavelengths and \(f\)-values from Morton (2003) with updates by Jenkins & Tripp (2006).

\(^b\) Rest frame equivalent width and error.

\(^c\) Blended line.
3.6 Chemical Composition of the DLA in front of UM 673A

Figure 3.7: A selection of metal absorption lines associated with the DLA in UM 673A, including transitions from neutral, singly and doubly ionised species. The data are shown as black histograms, while the red continuous lines are profile fits computed with VPFIT (see text). A light blue continuous line is used to indicate nearby absorption not associated with the DLA, such as the blue wing of Al\text{II} \lambda 1670 (top right-hand panel), which has been included in the profile fitting procedure. The y-axis is residual intensity, and the velocities on the x-axis are relative to the redshift of the absorption component with the highest optical depth (component number 3 in Table 3.3). Vertical tick marks indicate the velocities of the three absorption components producing the absorption lines, with the parameters listed in Table 3.3.
Table 3.3: Absorption components of low ion transitions in the DLA in line to UM 673A

<table>
<thead>
<tr>
<th>Comp. No.</th>
<th>( z_{\text{abs}} )</th>
<th>( \Delta v^a ) (km s(^{-1}))</th>
<th>( b ) (km s(^{-1}))</th>
<th>Fract.(^b)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>( 1.626348 \pm 2 \times 10^{-6} )</td>
<td>( -17.1 )</td>
<td>( 0.68 \pm 0.07 )</td>
<td>0.17</td>
</tr>
<tr>
<td>2</td>
<td>( 1.626454 \pm 2 \times 10^{-6} )</td>
<td>( -5.0 )</td>
<td>( 5.9 \pm 0.2 )</td>
<td>0.25</td>
</tr>
<tr>
<td>3</td>
<td>( 1.626498 \pm 8 \times 10^{-6} )</td>
<td>( 0.0 )</td>
<td>( 1.5 \pm 0.3 )</td>
<td>0.58</td>
</tr>
</tbody>
</table>

\(^a\)Velocity relative to \( z_{\text{abs}} = 1.626498 \)
\(^b\)Fraction of the total column density of Si II.

Table 3.4: Ion column densities of the DLA in UM 673A

<table>
<thead>
<tr>
<th>Ion</th>
<th>( \log N(X)/\text{cm}^{-2} ) C1(^a)</th>
<th>( \log N(X)/\text{cm}^{-2} ) C2+C3(^a)</th>
<th>( \log N(X)/\text{cm}^{-2} ) C1+C2+C3(^a)</th>
</tr>
</thead>
<tbody>
<tr>
<td>H I</td>
<td>N/A</td>
<td>N/A</td>
<td>20.7 ± 0.1</td>
</tr>
<tr>
<td>N I</td>
<td>( 12.17 \pm 0.41 )</td>
<td>( 14.97 \pm 0.25 )</td>
<td>( 14.97 \pm 0.25 )</td>
</tr>
<tr>
<td>Al II</td>
<td>( 11.91 \pm 0.16 )</td>
<td>( 12.89 \pm 0.07 )</td>
<td>( 12.93 \pm 0.08 )</td>
</tr>
<tr>
<td>Al III</td>
<td>( 11.05 \pm 0.16 )</td>
<td>( 11.89 \pm 0.03 )</td>
<td>( 11.95 \pm 0.05 )</td>
</tr>
<tr>
<td>Si II</td>
<td>( 13.99 \pm 0.07 )</td>
<td>( 14.67 \pm 0.02 )</td>
<td>( 14.75 \pm 0.03 )</td>
</tr>
<tr>
<td>Si III</td>
<td>( 12.32 \pm 0.43 )</td>
<td>( 13.05 \pm 0.09 )</td>
<td>( 13.12 \pm 0.17 )</td>
</tr>
<tr>
<td>S II</td>
<td>( 12.90 \pm 0.43 )</td>
<td>( 14.52 \pm 0.09 )</td>
<td>( 14.53 \pm 0.10 )</td>
</tr>
<tr>
<td>Ti II</td>
<td>N/A</td>
<td>N/A</td>
<td>( \leq 11.90^b )</td>
</tr>
<tr>
<td>Cr II</td>
<td>( 12.04 \pm 0.10 )</td>
<td>( 12.69 \pm 0.03 )</td>
<td>( 12.78 \pm 0.04 )</td>
</tr>
<tr>
<td>Fe II</td>
<td>( 14.05 \pm 0.05 )</td>
<td>( 14.44 \pm 0.02 )</td>
<td>( 14.59 \pm 0.03 )</td>
</tr>
<tr>
<td>Ni II</td>
<td>( 12.17 \pm 0.12 )</td>
<td>( 12.92 \pm 0.02 )</td>
<td>( 12.99 \pm 0.04 )</td>
</tr>
<tr>
<td>Zn II</td>
<td>( 10.54 \pm 0.43 )</td>
<td>( 11.37 \pm 0.09 )</td>
<td>( 11.43 \pm 0.15 )</td>
</tr>
</tbody>
</table>

\(^a\)C1/C2/C3: Component 1/2/3. \(^b\)3\(\sigma\) upper limit.

and C3. Although these components are unresolved with the HIRES instrumental resolution of FWHM = 8.3 km s\(^{-1}\) (Section 3.2), their narrow widths are indicated by the relative strengths and profiles of Si II lines of widely differing oscillator strengths. Ultimately, such low velocity dispersions can only be confirmed with higher resolution observations. However, I note that comparably low \( b \)-values are not unusual in cool clouds in the Milky Way disk and halo (e.g. Pettini, 1988; Barlow et al., 1995), and are now beginning to be measured at high redshifts too as the quality of the spectroscopic data improves (e.g. Jorgenson et al., 2009).

With the redshift \( z \) and velocity dispersion parameter \( b \) fixed to be the same for all lines of neutral and singly ionised species, I allow the column density in each component be the free parameter to be determined by \textsc{vpfit} [with the obvious restriction that all absorption lines arising from the same ground state of a given ion X D should yield the same value of \( N(\text{XD}) \)]. Values of \( N(\text{XD}) \) are collected in Table 3.4, where the column densities of components C2 and C3 are grouped together, as these two components are always blended at the resolution of the data. I also list in this Table the total (C1+C2+C3) column densities of each ion, which are better determined than those of the individual components.
Figure 3.8: A selection of weak metal absorption lines used for abundance determinations in the DLA in UM 673A, reproduced on an expanded scale. The data are shown as black histograms, while the red continuous lines are profile fits computed with VPFIT (see text). The $y$-axis is residual intensity, and the velocities on the $x$-axis are relative to the redshift of the absorption component with the highest optical depth (component number 3 in Table 3.3). Vertical red tick marks indicate the velocities of the three absorption components listed in Table 3.3.

It is important to stress in this regard that, with the exception of N I, the data include at least one weak, unsaturated, transition for all of the species used in the abundance determinations. Some examples are shown on an expanded scale in Figure 3.8. Under these circumstances, the values of column density deduced do not depend on the details of the profile fitting, because the lines lie on the linear part of the curve of growth. Although I record C II $\lambda$1334 and O I $\lambda$1302, these transitions are saturated; these species are therefore not included in Table 3.4 and in the subsequent abundance analysis. I do however include Ti II, whose $\lambda\lambda$1910.61, 1910.95 doublet is undetected; the 3$\sigma$ upper limit $\log [N$(Ti II)/cm$^{-2}] \leq 11.90$ in Table 3.4 was deduced assuming these lines to be as wide as Zn II $\lambda$2026, which is the weakest feature that I detect.

The column densities of the neutrals and first ions in Table 3.4 allow me to deduce directly the abundances of the corresponding elements. Before doing so in Section 3.6.2 below, I briefly comment on the absorption from more highly ionised gas.

Intermediate ionisation stages

This data cover two second ions, Al III $\lambda\lambda$1854, 1862 and Si III $\lambda$1206; all three transitions are shown in Figure 3.7. The weak Al III doublet lines are well reproduced by the same ‘cloud model’ determined for the low ionisation species (Table 3.3). The stronger Si III $\lambda$1206 line shows additional redshifted absorption which presumably arises in ionised gas, since it is absent from lines of comparable strength of ions which are dominant in H I regions (e.g. Si II $\lambda$1526—see Figure 3.7). Accordingly, the column densities of Al III and Si III listed in Table 3.4 refer only to the velocity interval appropriate to the neutral gas. These values of column densities
Figure 3.9: Transitions from highly ionised gas at redshifts close to that of the DLA in UM 673A. In each plot, the y-axis is residual intensity, and the velocities on the x-axis are relative to the redshift of the absorption component with the highest optical depth in neutral gas (component number 3 in Table 3.3). The data are shown as black histograms, while the red continuous lines are profile fits computed with the model parameters listed in Table 3.5. A light blue continuous line is used to indicate nearby absorption not associated with the DLA, but which has been included in the profile fitting procedure. The relative velocities of the three components of the model are indicated by red tick marks above the continuum level. All three components are blended with other absorption lines in Si IV λ1393 (top right-hand panel), but component number 3, at Δv = +114 km s⁻¹, is clearly absent in Si IV λ1402.

are useful for constraining the magnitude of putative ionisation corrections to the abundance determinations, as discussed in Section 3.6.2.

High ions

I also find lines from highly ionised gas at redshifts close to, but not the same as, that of the DLA. The extensive work by Fox et al. (2007) has shown this to be the case in many DLAs. The HIRES spectrum includes absorption lines from the C IV λλ1548, 1550 and Si IV λλ1393, 1402 doublets, although out of these four lines only C IV λ1550 is not blended and therefore affords the clearest view of the kinematic structure of the highly ionised gas (see Figure 3.9).

The highly ionised gas appears to be spread over three velocity components (Table 3.5); two are close in redshift to the DLA itself (Δv = −9.4 and +12.2 km s⁻¹ respectively for components 1 and 2 in Table 3.5), but the third is redshifted by Δv = +114 km s⁻¹. This third component must be of high ionisation indeed, as it is the strongest in C IV and yet is absent in Si IV (see Figure 3.9). Comparison of the b-values in Tables 3.5 and 3.3 shows that the high ions have larger velocity dispersions than the neutral gas. Table 3.6 lists the column densities of C IV and Si IV.

Comparing Figure 3.9 with the lower three panels of Figure 3.3, it can be readily appreciated that, in stark contrast with the low ion absorption lines, the C IV and Si IV absorption lines show
3.6 Chemical Composition of the DLA in front of UM 673A

Table 3.5: Absorption components of high ion transitions near the DLA in UM 673A

<table>
<thead>
<tr>
<th>Comp. No.</th>
<th>(z_{\text{abs}}) (km s(^{-1}))</th>
<th>(\Delta \nu^a) (km s(^{-1}))</th>
<th>(b) (km s(^{-1}))</th>
<th>Fract.(^b)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>1.626416 ± 3 \times 10^{-6}</td>
<td>-9.4</td>
<td>4.2 ± 0.9</td>
<td>0.51</td>
</tr>
<tr>
<td>2</td>
<td>1.626605 ± 3 \times 10^{-6}</td>
<td>+12.2</td>
<td>11.4 ± 0.7</td>
<td>0.11</td>
</tr>
<tr>
<td>3</td>
<td>1.627498 ± 2 \times 10^{-6}</td>
<td>+114.1</td>
<td>7.9 ± 0.3</td>
<td>0.38</td>
</tr>
</tbody>
</table>

\(^a\)Velocity relative to \(z_{\text{abs}} = 1.626498\)

\(^b\)Fraction of the total column density of C\(\text{IV}\).

Table 3.6: High ion column densities at redshifts close to that of the DLA in UM 673A

<table>
<thead>
<tr>
<th>Ion</th>
<th>(\log N(X)/\text{cm}^{-2}) (C1^a)</th>
<th>(\log N(X)/\text{cm}^{-2}) (C2^a)</th>
<th>(\log N(X)/\text{cm}^{-2}) (C3^a)</th>
</tr>
</thead>
<tbody>
<tr>
<td>C(\text{IV})</td>
<td>12.55 ± 0.06</td>
<td>13.10 ± 0.02</td>
<td>13.22 ± 0.01</td>
</tr>
<tr>
<td>Si(\text{IV})</td>
<td>12.59(^b)</td>
<td>12.45(^b)</td>
<td>(\leq 11.75^c)</td>
</tr>
</tbody>
</table>

\(^a\)C1/C2/C3: Component 1/2/3.

\(^b\)Uncertain because of blending.

\(^c\)3\(\sigma\) upper limit.

little variation between UM 673A and B. Applying the same VPFIT analysis as above to the C\(\text{IV}\) lines in UM 673B, returns values of redshift for the three absorption components which differ by less than 5 km s\(^{-1}\) from those listed in Table 3.5, and values of column density \(N(\text{C\(\text{IV}\)})\) which differ by less than a factor of 3 from those listed in Table 3.6. The finding that highly ionised gas has a much larger coherence scale than that of DLAs is not surprising, and in line with the results of earlier work on other QSO pairs (e.g. Rauch et al., 2001; Ellison et al., 2004, and references therein) and more recently on galaxy-galaxy pairs (Steidel et al., 2010).

3.6.2 Element Abundances

Abundance measurements for the DLA in UM 673A are collected in Table 3.7 and shown graphically in Figure 3.11. These values were deduced directly by dividing the column densities of ions which are dominant in H\(\text{I}\) regions by \(N(\text{H\(\text{I}\)})\) (see Table 3.4), with the implicit assumption that corrections for ionised gas and unseen ion stages are negligible so that, for example, \(N(\text{Si\(\text{II}\)})/N(\text{H\(\text{I}\)}) \equiv \text{Si}/\text{H}\).

This is normally a safe assumption at the high neutral hydrogen column densities of DLAs, where the gas is self-shielded from ionising radiation (e.g. Vladilo et al., 2001). However, given the unusual abundance pattern uncovered here, it is worthwhile reexamining the assumption that, for example, \(N(\text{N\(\text{I}\)})/N(\text{H\(\text{I}\)}) \equiv \text{N}/\text{H}\), and assess to what degree, if any, N may be over- or under-ionised compared to H (and the same for the other elements considered).
Table 3.7: ELEMENT ABUNDANCES IN THE DLA TOWARDS UM 673A

<table>
<thead>
<tr>
<th>Element</th>
<th>log (X/H)</th>
<th>[X/H]</th>
</tr>
</thead>
<tbody>
<tr>
<td>N</td>
<td>−4.17</td>
<td>−1.56 ± 0.25</td>
</tr>
<tr>
<td>Al</td>
<td>−5.56</td>
<td>−2.21 ± 0.08</td>
</tr>
<tr>
<td>Si</td>
<td>−4.49</td>
<td>−1.46 ± 0.03</td>
</tr>
<tr>
<td>S</td>
<td>−4.86</td>
<td>−1.31 ± 0.10</td>
</tr>
<tr>
<td>Ti</td>
<td>−7.09</td>
<td>&lt; −1.71</td>
</tr>
<tr>
<td>Cr</td>
<td>−6.36</td>
<td>−1.56 ± 0.04</td>
</tr>
<tr>
<td>Fe</td>
<td>−4.53</td>
<td>−1.58 ± 0.03</td>
</tr>
<tr>
<td>Ni</td>
<td>−5.79</td>
<td>−1.92 ± 0.04</td>
</tr>
<tr>
<td>Zn</td>
<td>−7.37</td>
<td>−1.90 ± 0.15</td>
</tr>
</tbody>
</table>

a Solar abundance scale reproduced from Appendix A.

To this end, I ran a suite of CLOUDY photoionisation models (Ferland et al., 1998), assuming that the DLA can be approximated by a slab of constant density gas in the range $−3 < \log\left[ n(\text{H})/\text{cm}^{-3} \right] < 3$. In these simulations, I included the Haardt & Madau (2001) metagalactic ionising background, as well as the cosmic microwave background, both at the redshift of the DLA. I adopted the solar abundance scale in Appendix A and globally scaled the metals to $Z_{\text{DLA}} = 1/30 Z_\odot$. No relative element depletions were employed, nor were grains added. The simulations were stopped when the column density of the DLA was reached. I am then able to calculate the ionisation correction, IC(X), for element X in ionisation stage N by the relation

$$IC(X) = \log \left[ \frac{N(X)}{N(\text{H})} \right]_{\text{intrinsic}} - \log \left[ \frac{N(XN)}{N(\text{HI})} \right]_{\text{computed}}$$

which will be negative when overestimating the abundance of an element by assuming that the dominant ionisation stage is representative of the true abundance.

The results of this exercise are shown in Figure 3.10, where it can be seen (left panel) that the ionisation corrections are less than 0.1 dex for most elements considered, for gas densities in excess of $\log\left[ n(\text{H})/\text{cm}^{-3} \right] = −2$. One can constrain the density by considering the relative column densities of successive ion stages, in this case using the Al III/Al II and Si III/Si II pairs. As can be seen from the right panel in Figure 3.10, both ratios give consistent answers, indicating a density $\log\left[ n(\text{H})/\text{cm}^{-3} \right] = −1.1 ± 0.1$ (from the better determined Si III/Si II ratio). At these densities, the ionisation corrections for most of the elements considered here are smaller (typically of the order $\sim 0.05$ dex) than the uncertainties in the corresponding ion’s column densities (shown by the height of the black boxes in Figure 3.11), except Al and S, whose true abundances may be higher than their entries in Table 3.7 by 0.1 dex and 0.2 dex respectively. In any case, it can be seen that ionisation corrections are a small effect, and can thus be safely neglected. The same conclusion is reached if the radiation field responsible for ionising the gas has a purely stellar origin, rather than the mix of Active Galactic Nuclei and star-forming galaxies that is the source of the metagalactic background considered by Haardt & Madau (2001).
3.6 Chemical Composition of the DLA in front of UM 673A

As an aside, one can also estimate the line-of-sight distance through the DLA, $D_{\text{los}}$, from the derived volume density under the assumption of constant density [i.e. $D_{\text{los}} = N(H_{\text{I}})/n(H)$]. I find $D_{\text{los}} = 2.1^{+1.0}_{-0.7}\,\text{kpc}$, consistent with the absence of the DLA in UM 673B and in agreement with the characteristic sizes deduced from the analysis presented in Section 3.4.

Returning now to Figure 3.11, the first conclusion to be drawn is that the $z_{\text{abs}} = 1.62650$ DLA in line to UM 673A is metal-poor, with an overall metallicity $Z \simeq 1/30\,Z_\odot$, or $-1.5$ on a log scale. Such low metallicity is not unexpected, given the narrow widths of the absorption lines (which are actually significantly narrower than expected on the basis of the relationship proposed by Prochaska et al., 2008) and the apparently low level of star formation deduced in Section 3.5.

When inspecting the relative abundances of the elements measured, however, I arrive at the interesting conclusion that this DLA exhibits some notable differences from the ‘average’ populations of DLAs and Galactic halo stars of similar metallicity, as I now point out. I consider two ‘average’ sets of DLAs, each assembled from the HIRES DLA database compiled by Prochaska et al. (2007). In each case I selected from the database all DLAs whose abundance is within a factor of 2 of a reference element, using in turn Si and Fe as the reference. Thus, the green vertical lines in the top panel of Figure 3.11 show the dispersion in the abundances of the elements considered here in all DLAs with $-1.76 \leq [\text{Si}/H] \leq -1.16$, i.e. within a factor of $\sim 2$ of the value $[\text{Si}/H] = -1.46 \pm 0.03$ I measure towards UM 673A. The middle panel shows the same data for all DLAs with $-1.86 \leq [\text{Fe}/H] \leq -1.26$. The numbers below the element tags in Figure 3.11 indicate the number of DLAs included in the average population for the corresponding element. If there exists only one measurement, the size of the error bar represents the uncertainty in that single measurement. If just two or three measurements exist, I instead plot the standard error in the mean. Otherwise, I use a robust determination of the dispersion in the average population (see Section 3.4).
Figure 3.11: All three panels show the abundances of the nine elements indicated at the top in the $z_{\text{abs}} = 1.62650$ DLA in front of UM 673A (black boxes); the size of each box corresponds to the uncertainty in each determination (except for Ti, for which I only have an upper limit). The green vertical lines in the top and middle panels show the dispersions of the abundances of each element in samples of DLAs that have, respectively, [Si/H] and [Fe/H] within a factor of 2 of the DLA in UM 673A. The numbers below the element tags indicate the number of DLA measurements available for each element (see text for further details). In the bottom panel the DLA chemical composition is compared with that of an ‘average’ population of Galactic metal-poor stars (open boxes) that have [Fe/H] within a factor of 2 of the DLA.
3.6 Chemical Composition of the DLA in front of UM 673A

Figure 3.12: Zn and Fe abundances in DLAs from the compilations by Prochaska et al. (2007), Dessauges-Zavadsky et al. (2007), and Noterdaeme et al. (2008) are shown with black circles, while the red dot is for the DLA in UM 673A. Note that the fall off in [Zn/Fe] with decreasing [Zn/H] in DLAs is not thought to be intrinsic, but rather to reflect a decreasing fraction of Fe in dust grains at lower overall metallicities, as measured by the normally undepleted Zn. The green shaded region shows the behaviour of [Zn/Fe] in Galactic stars (Saito et al., 2009, and references therein), with the width of the band corresponding to the ±1σ dispersion of values at a given metallicity. The dashed ellipse indicates the approximate locus of stars in the dwarf spheroidal galaxies Draco, Ursa Minor, and Sextans from the work by Shetrone, Côté, & Sargent (2001) and Cohen & Huang (2009).

Similarly, in the bottom panel of Figure 3.11, I overlay (empty boxes) the element abundances of the average population of Galactic metal poor stars drawn from the samples of Gratton et al. (2003) and Nissen et al. (2007), using Fe as the reference element. For all available measurements, the height of the box represents the dispersion in the average population.

Considering first the top panel of Figure 3.11, it appears that the DLA in UM 673A is Fe-rich and Zn-poor relative to other DLAs with similar Si abundance. There may be offsets in N and S too, but their statistics are poorer. The best DLA statistics are those for the middle panel, where it can be seen that, relative to other DLAs with similar Fe abundance, the absorber in UM 673A is deficient in Ti, Ni, and Zn. The same conclusion is reached by comparing with the stellar abundances in the bottom panel.

It seems unlikely that these apparently anomalous abundances are due to dust depletion, given that: (a) Ti, Ni, and Zn are depleted to very different degrees in Galactic dust (Savage & Sembach, 1996), and (b) depletions are in any case expected to introduce very minor corrections when the overall metallicity is as low as 1/30 of solar (e.g. Akerman et al., 2005).

The unusually low abundance of Zn (an element which is not readily incorporated into dust grains and thus normally provides a reliable benchmark for the abundance of Fe-peak elements, e.g. Pettini et al., 1990), can best be appreciated from Figure 3.12, where the value of [Zn/Fe] in UM 673A is compared with those measured in DLAs and Galactic stars. In the Figure, the
trend of decreasing $[\text{Zn}/\text{Fe}]$ with decreasing $[\text{Zn}/\text{H}]$ in DLAs is most naturally explained by the reduced dust depletion of Fe at lower metallicities, until solar relative abundance of the two elements is recovered when $[\text{Zn}/\text{H}] \approx -1.5$. In Galactic stars, Zn and Fe track each other closely over two decades in metallicity (from solar to $\sim 1/100$ solar), and Zn becomes progressively overabundant relative to Fe with decreasing metallicity when $[\text{Zn}/\text{H}] \lesssim -2$. Evidently, the $[\text{Zn}/\text{Fe}]$ ratio is lower in UM 673A than in any other DLA in current samples; similarly, none of the metal-poor stars in the compilations by Nissen et al. (2007) and Saito et al. (2009) exhibits an underabundance of Zn relative to Fe as pronounced as that uncovered here.

I have searched for clues in the composition of Galactic stars that are much more metal-poor than 1/30 of solar, in the extreme regime where chemical anomalies due to enrichment by only a few prior episodes of star formation may manifest themselves, but found none. The works by Cayrel et al. (2004) and Lai et al. (2008) show that when $[\text{Fe}/\text{H}] \ll -2$, Ti is overabundant by a factor of $\sim 2$ relative to Fe, Cr is progressively underabundant with decreasing $[\text{Fe}/\text{H}]$, Ni/Fe remains solar, and Zn exhibits the behaviour illustrated in Figure 3.12, quite unlike the relative abundances of these four elements relative to Fe in the DLA towards UM 673A.

Recently, data have become available on the abundances of these elements in individual stars of dwarf spheroidal galaxies in the Local Group (Shetrone, Côté, & Sargent, 2001; Cohen & Huang, 2009) and in ultra-faint dwarf companions of the Milky Way (Frebel et al., 2010). These measurements are of particular interest, as some of these galaxies may well have experienced different star formation histories from the stellar population(s) of the Galactic halo, and may therefore offer a different perspective for the interpretation of element ratios. It is thus intriguing to find that stars with $[\text{Fe}/\text{H}] \lesssim -2$ in the Draco, Sextans, and Ursa Minor dwarf spheroidal galaxies can exhibit sub-solar $[\text{Zn}/\text{Fe}]$ ratios (see Figure 3.12), and solar or sub-solar $[\text{Ti}/\text{Fe}]$ (in contrast with the super-solar $[\text{Ti}/\text{Fe}]$ of Galactic halo stars). However, the resemblance to the DLA in UM 673A does not extend to other elements: in the dwarf spheroidal stars observed by Shetrone, Côté, & Sargent (2001) and Cohen & Huang (2009) the mean value of $[\text{Ni}/\text{Fe}]$ is approximately solar and $[\text{Cr}/\text{Fe}]$ is mostly subsolar, whereas the opposite is found in the DLA (see Figure 3.11). Thus, while these data do offer examples of clear departures of some element ratios from the well-established Galactic halo pattern, a definite chemical correspondence between the DLA in UM 673A and nearby dwarf spheroidal galaxies cannot be established.

It is possible that, as more data of comparably high S/N ratio to those presented here are obtained for other very metal-poor DLAs, more examples of anomalous element abundances will be uncovered, challenging current calculations of stellar yields at low metallicities. Sub-solar $[\text{Zn,Ni,Ti}/\text{Fe}]$ ratios in very metal-poor DLAs may well be more common than has been appreciated so far, because the absorption lines of all three elements become vanishingly small when $[\text{Fe}/\text{H}] \lesssim -1.5$ (see, for example, Pettini et al., 2008a), requiring data of unusually high S/N ratio for their abundances to be determined.
3.7 Summary and Conclusions

Thanks to the high efficiency of the HIRES spectrograph at ultraviolet wavelengths, I have uncovered a damped Ly\(\alpha\) system at \(z = 1.62650\) in the spectrum of the gravitationally lensed QSO UM 673; the DLA had been overlooked until now, despite the many observations of this bright QSO over the last quarter of a century. From the analysis of high resolution and S/N ratio spectra of each image of the UM 673A,B pair, I draw the following conclusions.

(i) In the direction probed, the transverse extent of the DLA is much less than 2.7 kpc (this being the separation of the two sightlines at \(z = 1.62650\)), since I measure a drop by a factor of at least 400 in the column density of neutral hydrogen from UM 673A (log\([N(\text{H}I)/\text{cm}^{-2}]\) = 20.7) to UM 673B (log\([N(\text{H}I)/\text{cm}^{-2}]\) \(\leq 18.1\)). A comparable drop is seen in the column density of Si \(\text{II}\) and presumably other metals in stages that are dominant in H\(\text{I}\) regions.

(ii) From a reassessment of data on other QSO pairs in the literature, together with the new results for UM 673A,B, I find that, if the radial profile of \(N(\text{H}I)\) in DLAs declines exponentially, the median \(e\)-folding scale length is \(S_{\text{DLA},e} = 1.3 \pm 0.8\) kpc, smaller than had previously been realised. For a spherical cloud, this corresponds to a typical DLA radius \(R \approx 5 \pm 3\) kpc.

(iii) Towards UM 673B, I detect a weak and narrow Ly\(\alpha\) emission line at the same redshift as the DLA in UM 673A. If the line is due to recombination in H\(\text{II}\) regions, which I consider to be the most likely interpretation, its low luminosity \((L(\text{Ly}\alpha) = 4.3 \times 10^{41}\text{ erg s}^{-1})\) implies a modest star formation rate, \(SFR \approx 0.2\) M\(\odot\) yr\(^{-1}\). However, this is probably a lower limit considering: (a) the ease with which resonant Ly\(\alpha\) photons can be destroyed or scattered out of the line of sight, and (b) the limited spatial sampling of the narrow HIRES entrance slit.

(iv) In contrast with neutral gas, absorption by C \(\text{IV}\) exhibits only modest variations between the two sightlines. Evidently, highly ionised gas extends over much larger physical dimensions than the DLA, in accord with earlier conclusions from other QSO pairs and recent work on galaxy-galaxy pairs.

(v) The DLA in front of UM 673A is metal poor, with overall metallicity \(Z_{\text{DLA}} \approx 1/30Z_\odot\), and has a very simple velocity structure, with three absorption components spread over a narrow velocity interval, \(\Delta v = 17\) km s\(^{-1}\). Two of the components have very small internal velocity dispersions, with \(b = \sqrt{2}\sigma \approx 1\) km s\(^{-1}\), where \(\sigma\) is the one-dimensional rms velocity of the absorbing ions projected along the line of sight.

(vi) I am able to determine with precision the relative abundances of nine chemical elements, from N to Zn, thanks to the large number of metal absorption lines recorded at high S/N ratio. There appear to be some peculiarities in the detailed chemical make-up of the DLA, with the elements Ti, Ni, and Zn being deficient by factors of \(\sim 2–3\) compared to other DLAs and to Galactic halo stars of similar overall metallicity. The [Zn/Fe] ratio is the lowest measured in existing samples of DLAs and halo stars. While comparably low values of [Zn/Fe] have been
measured in some stars of nearby dwarf spheroidal galaxies, other element ratios differ between these stars and the DLA, so that a direct chemical correspondence cannot be established. An interpretation of these peculiar element ratios in terms of the previous history of chemical enrichment of the gas is still lacking.

Taken together, the small size, quiescent kinematics, and near-pristine chemical composition of the DLA in front of UM 673A would suggest an origin in a low-mass galaxy. However, the detection of nearby Ly$\alpha$ emission adds a new ‘twist’ to this picture. Although not intersected by this line of sight, there must presumably be a significant mass of cold gas within $\sim 3$ kpc of the DLA to fuel the star formation rate that I deduce from the Ly$\alpha$ luminosity. If this is the case, then the properties I measure may refer to an interstellar cloud—or complex of clouds—within a larger galaxy, rather than being representative of the whole galaxy. This caveat may also apply to other DLAs whose transverse dimensions have been probed with QSO pairs. However, the fact that in none of the cases studied so far has common DLA absorption been found over scales of a few kpc (see Table 3.1) points to one (or both) of two possibilities: either interstellar clouds with $N(\text{H} I) \geq 2 \times 10^{20}$ cm$^{-2}$ have covering fractions $f_c \ll 1$ within the interstellar media of high redshift galaxies or, if $f_c \sim 1$, the host galaxies of most DLAs are genuinely of small extent. The generally low metallicities of most DLAs independently point to an origin in low luminosity galaxies as discussed, among others, by Fynbo et al. (2008).

Looking ahead, the nature of the DLA studied here would undoubtedly be clarified by integral field observations of UM 673 at the wavelength of the H$\alpha$ emission line (which at $z = 1.62650$ is redshifted into the near-infrared $H$-band, at $\lambda_{\text{obs}} = 1.7241 \mu m$). Its detection would confirm the presence of a star-forming region, its spatial extent, and whether or not it is lensed by the foreground galaxy. As a closing remark, I also point out that, with its narrow velocity width and low metallicity, the DLA in UM 673A is a prime candidate for a rare measurement of the primordial D/H ratio (Pettini et al., 2008b). Such data, however, can only be obtained with the Hubble Space Telescope because the higher order Lyman lines in which the isotope shift could be resolved all lie at ultraviolet wavelengths which are inaccessible from the ground.
I present a high spectral resolution survey of the most metal-poor DLAs aimed at probing the nature and nucleosynthesis of the earliest generations of stars. This survey comprises 22 systems with iron abundance less than 1/100 solar; observations of seven of these are reported here for the first time. Together with recent measures of the abundances of C and O in Galactic metal-poor stars, I reinvestigate the trend of C/O in the very metal-poor regime and I compare, for the first time, the O/Fe ratios in the most metal-poor DLAs and in halo stars. I confirm the near-solar values of C/O in DLAs at the lowest metallicities probed, and find that their distribution is in agreement with that seen in Galactic halo stars. I find that the O/Fe ratio in very metal-poor (VMP) DLAs is essentially constant, and shows very little dispersion, with a mean $\langle \text{O/Fe} \rangle = +0.39 \pm 0.11$, in good agreement with the values measured in Galactic halo stars when the oxygen abundance is measured from the [O I] $\lambda 6300$ line. I speculate that such good agreement in the observed abundance trends points to a universal origin for these metals. In view of this agreement, I construct an abundance pattern for a typical very metal-poor DLA and compare it to model calculations of Population II and Population III nucleosynthesis to determine the origin of the metals in VMP DLAs. These results suggest that the most metal-poor DLAs may have been enriched by a generation of metal-free stars; however, given that abundance measurements are currently available for only a few elements, I am unable to rule out an additional contribution from Population II stars.
4.1 Introduction

The initial conditions for cosmic chemical evolution are of fundamental importance to our understanding of galaxy formation and the process of galactic chemical evolution. These conditions, set by the yields of the first few generations of stars, depend on various (largely unknown) factors including the form of the primordial stellar initial mass function and the uniformity of the enrichment of the intergalactic medium (IGM; Bromm & Larson 2004; Karlsson, Bromm, & Bland-Hawthorn 2011). In order to pin down the initial conditions of cosmic chemical evolution, one should seek to understand the origin and relative abundances of the metals in the least chemically evolved systems.

The most metal-poor DLAs, for example, are usually interpreted as distant protogalaxies at an early stage of chemical evolution (Erni et al., 2006). Whilst the origin of their metals is still largely unknown, recent hydrodynamical simulations suggest that such systems might have been enriched by just a few supernova events (Bland-Hawthorn et al., 2011). If this is indeed the case, the most metal-poor DLAs provide a simple route to study the first stages of chemical enrichment in our Universe.

By definition, DLAs have a neutral hydrogen column density in excess of $2 \times 10^{20}$ H I atoms cm$^{-2}$ (Wolfe et al. 1986; see also the review by Wolfe et al. 2005), which acts to self-shield the gas from the ultraviolet background radiation of quasars and galaxies (Haardt & Madau, 2001). This results in the gas having a simple ionisation structure subject to negligible corrections for unseen ion stages (Vladilo et al., 2001), quite unlike the Ly$\alpha$ forest clouds that trace the low density regions of the IGM (e.g. Simcoe, Sargent, & Rauch 2004). The main concerns that limit abundance studies in DLAs are line saturation and the possibility that dust may hide some fraction of the metals (Vladilo, 2004). These concerns are alleviated when the metallicity of the DLA is below $\sim 10^{-2} Z_{\odot}$, which is also the regime where one expects to uncover the enrichment signature of the earliest generations of stars.

The recent interest in the most metal-poor DLAs (Pettini et al., 2008a; Penprase et al., 2010) complements the ongoing local studies of metal-poor stars in the halo of the Milky Way (Cayrel et al., 2004; Beers & Christlieb, 2005; Suda et al., 2008; Frebel et al., 2010). These stars are believed to have condensed out of near-pristine gas (perhaps a metal-poor DLA itself?), that was enriched by only a few earlier generations of stars. Thus, the first generation of stars can also be studied through the signature retained in the stellar atmospheres of the most metal-poor stars in the halo of our Galaxy. However, unlike the relative ease with which one can measure the abundances of metal-poor DLAs, deriving element abundances from the stellar atmospheres of metal-poor stars is not straightforward (Asplund, 2005). Systematic uncertainties in the derived abundances are introduced by assuming that the spectral line being examined forms in a region that is in local thermodynamic equilibrium (LTE), as well as the need to account for three-dimensional (3D) effects in the 1D stellar atmosphere models.

These effects are particularly acute for oxygen, where several different abundance indicators
are known to produce contradictory estimates in the low-metallicity regime (García Pérez et al., 2006). Despite the efforts of many authors, our uncertainty in the derived oxygen abundances has sparked an ongoing debate as to the trend of [O/Fe] in the Milky Way when [Fe/H] \( \lesssim -1.0 \). A history of the relevant discussion on [O/Fe] is provided by McWilliam (1997), with further details given in Section 4.6.2. In brief, at low metallicity, both O and Fe are produced exclusively by type-II supernovae (SNe II) and the winds from their progenitors. When [Fe/H] \( \lesssim -1.0 \), there is a drop in [O/Fe] due to the delayed contribution of Fe from type-Ia supernovae (SNe Ia). Thus, the [O/Fe] ratio is most commonly used to measure the time delay between SNe II and the onset of SNe Ia. At the lowest metallicity, however, one can use the [O/Fe] ratio as a measure of the relative production of \( \alpha \)- to Fe-peak elements by the first few generations of massive stars.

Another key diagnostic ratio at low metallicity that may shed light on the nature of the early generations of stars was uncovered by Akerman et al. (2004) who reported a rather surprising evolution of [C/O] with decreasing O abundance in their sample of 34 halo stars (see also Spite et al. 2005). In disc and halo stars when the oxygen abundance is \( \gtrsim -1.0 \), [C/O] steadily rises from [C/O] \( \sim -0.5 \) to solar. When [O/H] \( \lesssim -1.0 \), galactic chemical evolution models that only consider the nucleosynthetic products of Population II stars predict [C/O] to decrease or plateau, contrary to the observed trend. The increase in [C/O] with decreasing metallicity has thus been interpreted as evidence for an increased carbon yield from either Population III stars (Chieffi & Limongi, 2002; Umeda & Nomoto, 2003; Heger & Woosley, 2010) or rapidly-rotating low-metallicity Population II stars (Chiappini et al., 2006). At first, concerns were raised regarding the accuracy of the derived C and O abundances, since the lines used are subject to large non-LTE corrections. Fabbian et al. (2009a), however, performed a non-LTE analysis of the same lines, with further contraints from additional C I lines, to confirm the reality of the stellar [C/O] trend. These results depend somewhat on the adopted cross sections for collisions of C I and O I atoms with electrons and hydrogen atoms, but for all probable values, [C/O] increases with decreasing metallicity when [O/H] \( < -2.0 \).

To summarise, at present there are still some remaining concerns that prevent one from accurately measuring C and O abundances in the atmospheres of metal-poor halo stars. These difficulties have prompted a few teams to focus on very metal-poor\(^1\) (VMP) DLAs where the absorption lines of C II and O I may be unsaturated and the abundances of C and O can be measured with confidence. Unfortunately, these near-pristine DLAs are rare, falling in the tail of the metallicity distribution function of DLAs (Prochaska et al., 2007). Thus, only a handful of confirmed VMP DLAs are known at present. The first high spectral resolution survey (\( R \approx 40000 \), full width at half maximum, FWHM \( \approx 7 \) km s\(^{-1}\)) for VMP DLAs was conducted by Pettini et al. (2008a), whose specific goal was to study the relative abundances of the CNO group of elements as a probe of early nucleosynthesis. Indeed, this was the first study to independently

\(^1\)Herein, I adopt “very metal-poor” to be those DLAs with [Fe/H] \( < -2.0 \), in line with the classification scheme for stars proposed by Beers & Christlieb (2005).
confirm the increased [C/O] abundance at low metallicity, suggesting that near-solar values of [C/O] are commonplace in this metallicity regime.

The [C/O] trend reported by Pettini et al. (2008a) has recently been confirmed by Penprase et al. (2010) in a medium spectral resolution \((R \approx 5000, \text{FWHM} \approx 60 \text{ km s}^{-1})\) survey of 35 DLAs. In many of their systems, the C II and O I lines were thought to be affected by line saturation, leaving only five DLAs to test the trend in C/O. Interestingly, this sample of DLAs suggests that [C/O] continues to rise to supersolar values when [O/H] \(\lesssim -3\).

Such surveys for VMP DLAs are most useful for studying the general properties of entire clouds of near-pristine gas before they form stars. In this contribution, I build on the ongoing survey for the most metal-poor DLAs as probes of early nucleosynthesis conducted by Pettini et al. (2008a). With additional systems drawn from the literature, the total sample presented herein amounts to 22 VMP DLAs with abundance measurements derived from high spectral resolution data. From this sample, I confirm the elevated [C/O] values in these systems and, for the first time, present the trend of [O/Fe] in the most metal-poor DLAs. For both of these diagnostic ratios, I comment on the implications these findings have for local studies of Galactic metal-poor halo stars. Finally, I construct the abundance pattern of a typical VMP DLA for the elements C, N, O, Al, Si and Fe, and compare it to model calculations of Population II and Population III nucleosynthesis.

This Chapter is arranged as follows. In Section 4.2 I detail the processing and preparation of the data. In Section 4.3 I explain the profile fitting procedure used for this new sample of VMP DLAs, which are discussed in Section 4.4. The accuracy of the abundance analysis is discussed in Section 4.5, before investigating the behaviour of [C/O] and [O/Fe] in VMP DLAs and comparing with stellar data, in Section 4.6. Finally, I discuss the implications for these findings in Section 4.7, before summarising the main results and drawing my conclusions in Section 4.8.

4.2 Observations and Data Reduction

4.2.1 Target Selection

Even at the relatively low spectral resolution afforded by the Sloan Digital Sky Survey (SDSS), one can easily recognise DLAs in the spectra of quasars, owing to the characteristic damping wings of the Ly\(\alpha\) absorption line profile. Subsequent identification of associated metal line absorption leads to a rough estimate of the gas-phase metallicity. Candidate metal-poor DLAs are then identified as those DLAs that appear to exhibit no metal line absorption; their absorption features are unresolved at the spectral resolution of the SDSS. However, when these candidates are re-observed with echelle spectrographs of high resolution \((R \gtrsim 30000, \text{FWHM} \lesssim 10 \text{ km s}^{-1})\), the metal absorption lines are resolved, and in many cases it is possible to measure elemental abundances with confidence (Pettini et al., 2008a).
4.2 Observations and Data Reduction

The most recent trawls through SDSS spectra of \( \sim 8000 \) quasars with \( z_{\text{em}} \gtrsim 2.2 \) has yielded a sample of \( \sim 1000 \) DLAs (Noterdaeme et al., 2009; Prochaska & Wolfe, 2009), of which \( \sim 400 \) are classified as ‘metal-poor’ (Penprase et al., 2010).\(^2\) From compilations such as these, I selected a handful of metal-poor DLA candidates that exhibit no discernible metal-line absorption at the spectral resolution of the SDSS, and re-observed these with echelle spectrographs, giving higher priority to candidates with: (i) bright quasars, so as to efficiently obtain spectra with signal-to-noise ratios \( S/N \gtrsim 20 \) in the continuum; (ii) DLAs where the difference between \( z_{\text{abs}} \) and \( z_{\text{em}} \) is minimised so that the absorption lines of interest (e.g. \( \text{O I}\lambda 1302 \) and \( \text{C II}\lambda 1334 \)) are not blended with unrelated Ly\( \alpha \) forest lines; (iii) quasars whose emission redshift is below \( z_{\text{em}} \lesssim 3.3 \), so there is an improved chance that other lines of interest (e.g. \( \text{O I}\lambda 1039 \) and \( \text{C II}\lambda 1036 \)) are not blended with Ly\( \alpha \) forest lines; (iv) DLAs at the low end of the column density distribution function – that are still DLAs – to ensure that even the strongest metal absorption lines are unsaturated, allowing me to measure the metal ion column densities with confidence; and (v) quasars with more than one metal-poor DLA candidate in their spectra.

An example SDSS spectrum that is known to contain a VMP DLA is shown in Figure 4.1. This survey to date consists of 12 DLAs with \([\text{Fe/H}] \leq -2.0\). Initial results for four of these were published by Pettini et al. (2008a), while a fifth DLA, showing a pronounced C enhancement relative to Fe, will be the subject of Chapter 5. The observations and analysis of the remaining seven DLAs, including one from the European Southern Observatory’s (ESO) Ultraviolet and Visual Echelle Spectrograph (UVES) data archive, are presented here. To these data, I add a collection of published abundance measurements in ten VMP DLAs, selected as described in Section 4.4.8, to assemble an overall sample of abundance measurements in 22 VMP DLAs, all obtained from high resolution spectra \( (R \gtrsim 30000) \).

4.2.2 Echelle spectroscopic follow-up

In order to achieve the high signal-to-noise ratio and spectral resolution required for accurate DLA abundance measurements, I observed the prime candidates with echelle spectrographs on 8–10 m class telescopes. Most of these candidates were observed with the UVES spectrograph (Dekker et al., 2000), which is mounted on UT2 at the Very Large Telescope facility. An additional system was observed with the W. M. Keck Observatory’s High Resolution Echelle Spectrometer (HIRES, Vogt et al. 1994) on the Keck I telescope. Table 4.1 lists details of the observations of seven VMP DLAs reported here for the first time. For J1340+1106, I include details of some additional data, of comparable spectral resolution to mine, retrieved from the ESO\(^3\) and Keck Observatory\(^4\) data archives (program IDs 67.A-0078(A) and U11H respectively). I have also retrieved UVES spectra of the quasar J0311–1722 from the ESO data

---

\(^2\)In this context, a DLA is classed as ‘metal-poor’ if it has fewer than three significantly (4\(\sigma\)) detected metal absorption lines at the spectral resolution of the SDSS.

\(^3\)http://archive.eso.org/eso/eso_archive_main.html

\(^4\)https://www2.keck.hawaii.edu/koa/public/koa.php
Figure 4.1: An example SDSS spectrum (shown by the black histogram) that contains a VMP DLA. The top panel presents the entire SDSS spectrum for the quasar J1001 + 0343, which was identified as a candidate VMP DLA. The remaining panels plot expanded versions of the strongest lines coincident with the DLA absorption. Even with this high signal-to-noise spectrum, there are no discernable metal lines.
Table 4.1: JOURNAL OF OBSERVATIONS FOR THE VERY METAL-POOR DLA SURVEY

<table>
<thead>
<tr>
<th>QSO</th>
<th>$g^a$ (mag)</th>
<th>$z_{em}$</th>
<th>$z_{abs}$</th>
<th>Telescope/ Instrument</th>
<th>Wavelength Range (Å)</th>
<th>Resolution (km s$^{-1}$)</th>
<th>Integration Time (s)</th>
<th>S/N$^b$</th>
<th>Program ID</th>
</tr>
</thead>
<tbody>
<tr>
<td>J0311−1722</td>
<td>17.7</td>
<td>4.039</td>
<td>3.73400</td>
<td>VLT/UVES</td>
<td>4370−6410$^c$</td>
<td>6.9</td>
<td>3600</td>
<td>15</td>
<td>69.A-0613(A)$^d$</td>
</tr>
<tr>
<td>J0831+3358</td>
<td>19.5</td>
<td>2.427</td>
<td>2.30364</td>
<td>KECK/HIRESb</td>
<td>3130−5970$^c$</td>
<td>7.3</td>
<td>18600</td>
<td>15</td>
<td>A185Hb</td>
</tr>
<tr>
<td>J1001+0343</td>
<td>19.2</td>
<td>3.198</td>
<td>3.07841</td>
<td>VLT/UVES</td>
<td>3740−6650$^c$</td>
<td>7.3</td>
<td>33700</td>
<td>40</td>
<td>083.A-0042(A)</td>
</tr>
<tr>
<td>J1037+0139</td>
<td>19.4</td>
<td>3.059</td>
<td>2.70487</td>
<td>VLT/UVES</td>
<td>3640−6650$^c$</td>
<td>7.3</td>
<td>26075</td>
<td>45</td>
<td>083.A-0042(A)</td>
</tr>
<tr>
<td>J1340+1106</td>
<td>19.0</td>
<td>2.914</td>
<td>2.50792</td>
<td>VLT/UVES &amp; KECK/HIRES</td>
<td>3483−6652$^c$</td>
<td>7.3</td>
<td>29800 &amp; 50</td>
<td>40</td>
<td>085.A-0109(A) &amp;</td>
</tr>
<tr>
<td></td>
<td></td>
<td></td>
<td>2.79583</td>
<td></td>
<td>3483−9396$^c$</td>
<td>10.3</td>
<td>10800 &amp; 40</td>
<td>25</td>
<td>67.A-0078(A)$^d$ &amp;</td>
</tr>
<tr>
<td>J1419+0829</td>
<td>18.9</td>
<td>3.030</td>
<td>3.04973</td>
<td>VLT/UVES</td>
<td>3710−6652$^c$</td>
<td>7.3</td>
<td>29800</td>
<td>43</td>
<td>085.A-0109(A)</td>
</tr>
</tbody>
</table>

$^a$Magnitudes are SDSS $g$-band, except for J0311−1722 (not covered by the SDSS) which is $R$-band (Péroux et al., 2001).

$^b$Indicative signal-to-noise ratio at 5000 Å (or 6000 Å in the case of J1340+1106).

$^c$With some wavelength gaps.

$^d$Spectra downloaded from either the UVES or HIRES data archives$^{4,5}$.
Chapter 4. A survey for the most metal-poor DLAs

archive (program ID 69.A-0613(A); see Péroux et al. 2005), since the metal lines for the VMP DLA along this sightline have not been previously analysed.

The UVES observations reported here [program IDs 083.A-0042(A) & 085.A-0109(A)] employed a 1.2″ wide slit, resulting in a spectral resolution $R \sim 41000$ (velocity FWHM $\approx 7.3$ km s$^{-1}$) sampled with $\sim 3$ pixels. Dichroic #1 was used to split the quasar light into the blue and red spectroscopic arms containing the HER$_5$ filter and SHP700 filter respectively. The resulting central wavelength for each arm was 3900 Å (blue) and 5640 Å (red). Both the blue- and red-sensitive CCDs used $2 \times 2$ on-chip binning. For the HIRES observations (program ID A185Hb) I used the C5 decker (a 7.0 × 1.148 arcsec slit) which, with sub-arcsec seeing, gave a spectral resolution of $R \sim 41000$ (cf. Chapter 5), also sampled with $\sim 3$ pixels. I employed the ultraviolet cross-disperser with no filters, and used $2 \times 2$ on-chip binning.

4.2.3 Data Reduction

I used the standard UVES data reduction pipeline\(^5\) provided by ESO to reduce the UVES data. The UVES reduction pipeline performs the usual steps relevant to echelle data reduction. The preliminary steps include bias subtraction, flat fielding, and background subtraction. The echelle orders are then traced using a flat field frame taken with a pinhole decker, and 1D spectra extracted. The data are wavelength calibrated with reference to a ThAr lamp.

The HIRES data were reduced with the MAKEE data reduction pipeline developed by Tom Barlow. MAKEE performs the same reduction steps as outlined above, but a trace frame is not always readily available. When available, the orders were traced using a flat-field frame taken with a pinhole decker. Otherwise, the science exposure of the quasar itself was used when a satisfactory trace could be made. Failing this, a trace frame was generated with a suite of purpose-built PYTHON programs, using the science frame as a guide to trace the echelle orders.

Following these initial reduction steps, for each object I combined the science exposures using the software package UVES_POPLER\(^6\), maintained by Michael Murphy. This software merges individual echelle orders, and maps the data onto a vacuum heliocentric wavelength scale. Finally, the data were normalised by dividing out the quasar continuum and emission lines. Using the approximate redshift derived from each DLA’s SDSS discovery spectrum, I then prepared the final data for analysis by extracting a ±150 km s$^{-1}$ window around the pixel with highest optical depth near all available absorption lines of interest. Finally, a further fine adjustment to the continuum was applied to these extracted portions of the spectra when necessary.

\(^{5}\)I used version 4.3.0, available from: http://www.eso.org/sci/software/pipelines/

\(^{6}\)UVES_POPLER can be downloaded from http://astronomy.swin.edu.au/~mmurphy/UVES_popler
Table 4.2: Absorption Components of Low Ion Transitions

<table>
<thead>
<tr>
<th>Component Number</th>
<th>Component</th>
<th>$z_{\text{abs}}$</th>
<th>$b$ (km s$^{-1}$)</th>
<th>Fraction$^a$</th>
<th>$N$(Si II)</th>
</tr>
</thead>
<tbody>
<tr>
<td>J0311$-$1722: DLAs at $z_{\text{abs}} = 3.73400$</td>
<td>1</td>
<td>3.733862 ± 0.000017</td>
<td>5.6 ± 1.2</td>
<td>0.18</td>
<td></td>
</tr>
<tr>
<td></td>
<td>2</td>
<td>3.733998 ± 0.000007</td>
<td>2.7 ± 0.8</td>
<td>0.23</td>
<td></td>
</tr>
<tr>
<td></td>
<td>3</td>
<td>3.734035 ± 0.000028</td>
<td>14.4 ± 1.5</td>
<td>0.20</td>
<td></td>
</tr>
<tr>
<td></td>
<td>4</td>
<td>3.734439 ± 0.000002</td>
<td>4.6 ± 0.2</td>
<td>0.39</td>
<td></td>
</tr>
<tr>
<td>J0831$+$3358: DLAs at $z_{\text{abs}} = 2.30364$</td>
<td>1</td>
<td>2.303565 ± 0.000004</td>
<td>4.4 ± 0.2</td>
<td>0.49</td>
<td></td>
</tr>
<tr>
<td></td>
<td>2</td>
<td>2.303720 ± 0.000004</td>
<td>6.1 ± 0.3</td>
<td>0.51</td>
<td></td>
</tr>
<tr>
<td>J1001$+$0343: DLAs at $z_{\text{abs}} = 3.07841$</td>
<td>1</td>
<td>3.078413 ± 0.000002</td>
<td>7.0 ± 0.1</td>
<td>1.00</td>
<td></td>
</tr>
<tr>
<td>J1037$+$0139: DLAs at $z_{\text{abs}} = 2.70487$</td>
<td>1</td>
<td>2.704870 ± 0.000002</td>
<td>5.9 ± 0.2</td>
<td>1.00</td>
<td></td>
</tr>
<tr>
<td>J1340$+$1106: DLAs at $z_{\text{abs}} = 2.50792$</td>
<td>1</td>
<td>2.507649 ± 0.000003</td>
<td>2.0 ± 0.4</td>
<td>0.19</td>
<td></td>
</tr>
<tr>
<td></td>
<td>2</td>
<td>2.507921 ± 0.000001</td>
<td>5.8 ± 0.1</td>
<td>0.81</td>
<td></td>
</tr>
<tr>
<td>J1340$+$1106: DLAs at $z_{\text{abs}} = 2.79583$</td>
<td>1</td>
<td>2.7955473 ± 0.0000019</td>
<td>9.3 ± 0.2</td>
<td>0.19</td>
<td></td>
</tr>
<tr>
<td></td>
<td>2</td>
<td>2.7958285 ± 0.0000005</td>
<td>6.45 ± 0.05</td>
<td>0.81</td>
<td></td>
</tr>
<tr>
<td>J1419$+$0829: DLAs at $z_{\text{abs}} = 3.04973$</td>
<td>1</td>
<td>3.049649 ± 0.000002</td>
<td>3.5 ± 0.1</td>
<td>0.43</td>
<td></td>
</tr>
<tr>
<td></td>
<td>2</td>
<td>3.049835 ± 0.000002</td>
<td>6.4 ± 0.1</td>
<td>0.57</td>
<td></td>
</tr>
</tbody>
</table>

$^a$Fraction of the total column density of Si II.

4.3 Profile Fitting

For DLAs with a metallicity below 1/100 $Z_{\odot}$, the metal line absorption is typically concentrated in only a few clouds of low velocity dispersion (Ledoux et al., 2006; Murphy et al., 2007; Prochaska et al., 2008). By assuming that a Maxwellian distribution accurately describes the velocities of the dominant atoms within the neutral cloud, one can model a DLA’s absorption lines by a Voigt profile. To this end, I employed the Voigt profile fitting software VPFIT to derive the cloud parameters for all DLAs in this sample.\textsuperscript{7}

VPFIT uses a chi-squared minimisation algorithm to simultaneously fit multiple Voigt profiles to a set of absorption lines characterised by three free parameters: (1) the cloud’s absorption redshift ($z_{\text{abs}}$); (2) the Doppler parameter of the absorbing gas ($b$ in km s$^{-1}$); and (3) the column density of the ion that gives rise to the absorption line. When it was evident that the DLA metal absorption arises from more than one cloud component, I introduced additional components to reduce the $\chi^2$, whilst maintaining realistic errors on the derived parameters (i.e. $\lesssim 10\%$ uncertainty on $b$ and a redshift uncertainty less than the sampling size of $\sim 2.5$ km s$^{-1}$). Throughout the fitting procedure I assumed that the dominant ions in H I regions (e.g. C II, N I, O I, Si II,\textsuperscript{7} VPFIT is available from http://www.ast.cam.ac.uk/~/rfc/vpfit.html)
Table 4.3: The Adopted Metal Line Laboratory Wavelengths and Oscillator Strengths

<table>
<thead>
<tr>
<th>Ion</th>
<th>Wavelength (Å)</th>
<th>f</th>
<th>Ion</th>
<th>Wavelength (Å)</th>
<th>f</th>
<th>Ion</th>
<th>Wavelength (Å)</th>
<th>f</th>
</tr>
</thead>
<tbody>
<tr>
<td>C II</td>
<td>1036.3367</td>
<td>0.118</td>
<td>Al II</td>
<td>1670.7886</td>
<td>1.740</td>
<td>Fe II</td>
<td>1063.1764</td>
<td>0.0547</td>
</tr>
<tr>
<td>C II</td>
<td>1334.5323</td>
<td>0.1278</td>
<td>Al III</td>
<td>1854.71829</td>
<td>0.559</td>
<td>Fe II</td>
<td>1081.8748</td>
<td>0.0126</td>
</tr>
<tr>
<td>C II*</td>
<td>1335.6627</td>
<td>0.01277</td>
<td>Al III</td>
<td>1862.79113</td>
<td>0.278</td>
<td>Fe II</td>
<td>1096.8769</td>
<td>0.0327</td>
</tr>
<tr>
<td>C II*</td>
<td>1335.7077</td>
<td>0.115</td>
<td>Si II</td>
<td>989.8731</td>
<td>0.171</td>
<td>Fe II</td>
<td>1125.4477</td>
<td>0.0156</td>
</tr>
<tr>
<td>N I</td>
<td>1134.4149</td>
<td>0.0278</td>
<td>Si II</td>
<td>1020.6989</td>
<td>0.0168</td>
<td>Fe II</td>
<td>1143.2260</td>
<td>0.0192</td>
</tr>
<tr>
<td>N I</td>
<td>1134.9803</td>
<td>0.0416</td>
<td>Si II</td>
<td>1190.4158</td>
<td>0.292</td>
<td>Fe II</td>
<td>1144.9379</td>
<td>0.0830</td>
</tr>
<tr>
<td>N I</td>
<td>1199.5496</td>
<td>0.1320</td>
<td>Si II</td>
<td>1193.2897</td>
<td>0.582</td>
<td>Fe II</td>
<td>1260.533</td>
<td>0.0240</td>
</tr>
<tr>
<td>N I</td>
<td>1200.2233</td>
<td>0.0869</td>
<td>Si II</td>
<td>1260.4221</td>
<td>1.18</td>
<td>Fe II</td>
<td>1608.4509</td>
<td>0.0577</td>
</tr>
<tr>
<td>N I</td>
<td>1200.7098</td>
<td>0.0432</td>
<td>Si II</td>
<td>1304.3702</td>
<td>0.0863</td>
<td>Fe II</td>
<td>1611.20034</td>
<td>0.00138</td>
</tr>
<tr>
<td>N II</td>
<td>1083.9937</td>
<td>0.111</td>
<td>Si II</td>
<td>1526.7070</td>
<td>0.133</td>
<td>Fe II</td>
<td>2344.21296</td>
<td>0.1142</td>
</tr>
<tr>
<td>O I</td>
<td>925.446</td>
<td>0.000354</td>
<td>Si II</td>
<td>1808.01288</td>
<td>0.00208</td>
<td>Fe II</td>
<td>2374.46033</td>
<td>0.0313</td>
</tr>
<tr>
<td>O I</td>
<td>936.6295</td>
<td>0.00365</td>
<td>Si II</td>
<td>1250.578</td>
<td>0.00543</td>
<td>Fe II</td>
<td>2382.76418</td>
<td>0.320</td>
</tr>
<tr>
<td>O I</td>
<td>948.6855</td>
<td>0.00631</td>
<td>Si II</td>
<td>1253.805</td>
<td>0.0109</td>
<td>Ni II</td>
<td>1317.217</td>
<td>0.057</td>
</tr>
<tr>
<td>O I</td>
<td>976.4481</td>
<td>0.00331</td>
<td>Si II</td>
<td>1259.5180</td>
<td>0.0166</td>
<td>Ni II</td>
<td>1370.132</td>
<td>0.056</td>
</tr>
<tr>
<td>O I</td>
<td>988.5778</td>
<td>0.000553</td>
<td>Ar I</td>
<td>1048.2199</td>
<td>0.263</td>
<td>Ni II</td>
<td>1454.842</td>
<td>0.0323</td>
</tr>
<tr>
<td>O I</td>
<td>988.6549</td>
<td>0.0083</td>
<td>Ar I</td>
<td>1066.6598</td>
<td>0.0675</td>
<td>Ni II</td>
<td>1709.6042</td>
<td>0.0324</td>
</tr>
<tr>
<td>O I</td>
<td>988.7734</td>
<td>0.0465</td>
<td>Cr II</td>
<td>2056.25693</td>
<td>0.1030</td>
<td>Ni II</td>
<td>1741.5531</td>
<td>0.0427</td>
</tr>
<tr>
<td>O I</td>
<td>1039.2304</td>
<td>0.00907</td>
<td>Cr II</td>
<td>2062.23610</td>
<td>0.0759</td>
<td>Ni II</td>
<td>1751.9157</td>
<td>0.0277</td>
</tr>
<tr>
<td>O I</td>
<td>1302.1685</td>
<td>0.048</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
</tbody>
</table>
4.4 Individual Objects

Fe II) are kinematically associated with the same gas. I therefore fixed the redshift and Doppler parameter of each absorption component to be the same for each of these ions. The resulting cloud model parameters are given in Table 4.2 where, as a guide, the last column lists the fraction of the total column density of Si II in each component. The total column densities of available ions in each DLA are collected in Appendix B. Table 4.3 lists laboratory wavelengths and oscillator strengths of relevant atomic transitions from the compilation by Morton (2003), with subsequent updates by Jenkins & Tripp (2006).

4.4 Individual Objects

In this section, I briefly comment on the properties of each new DLA analysed in this Chapter.

4.4.1 J0311−1722: DLA at $z_{\text{abs}} = 3.73400$

The VMP DLA along the line-of-sight to J0311−1722 (J2000.0: 03h11m15s, −17°22′47″4) was first identified by Péroux et al. (2001). Follow-up UVES spectroscopy by Péroux et al. (2005) revealed a Lyα absorber at $z_{\text{abs}} = 3.734$ with log [$N(\text{HI})$/cm$^{-2}$] = 19.48±0.10, which is lower than the conventional limit for DLAs set by Wolfe et al. (1986), log [$N(\text{HI})$/cm$^{-2}$] ≥ 20.3. Such systems are often referred to as sub-DLAs (Péroux et al. 2003a) or super Lyman-limit systems, and are defined to have 19.0 ≤ log [$N(\text{HI})$/cm$^{-2}$] ≤ 20.3.

Péroux et al. (2005) derived their estimate of $N(\text{HI})$ for this absorber from a consistent model fit to the first 6 Lyman lines (from Lyα to Ly6), where the Lyα line produces the poorest fit. Having re-analysed these data, giving higher priority to the Lyα and Lyβ lines, I find that a better fit results if log [$N(\text{HI})$/cm$^{-2}$] = 20.30±0.06 (see top panel of Figure 4.2), provided that the Doppler parameter is $b < \sim 30$ km s$^{-1}$. Possibly, the value of $N(\text{HI})$ derived by Péroux et al. (2005) was driven by the higher order Lyman lines, which may be blended with other Lyman forest lines, and resulted in a Doppler parameter much larger than 30 km s$^{-1}$.

In the lower panels of Figure 4.2 I present a selection of the metal lines associated with this VMP DLA. All of the available metal absorption lines are unsaturated, leading to reliable estimates of the metal column densities. A four component model was found to accurately reproduce the metal-line profiles, whilst maintaining reasonable estimates of the parameter errors (see Section 4.3). The derived cloud model parameters are listed in Table 4.2. Although both C II lines ($\lambda$1334 and $\lambda$1036) are blended on the red wing of component 4 (centred near $v = 0$ km s$^{-1}$ in Figure 4.2), this does not greatly affect the final estimate of $N(\text{C II})$, since component 4 has a well-determined Doppler parameter from the relative strengths of Si II $\lambda$1260 and Si II $\lambda$1304, and from other unblended transitions in regions of high signal-to-noise. In any case, the majority of the absorbing column is contributed by the first three components ($\sim 60\%$). I list the total column density returned by VPFit for each available ion in Table 4.4.
Figure 4.2: The top panel shows the DLA towards J0311−1722 (black histogram) which exhibits a damped Lyα line at $z_{\text{abs}} = 3.73400$. The red continuous line shows the theoretical Voigt profile for a neutral hydrogen column density $\log[\text{N}($H\text{I}$)/\text{cm}^{-2}] = 20.30$. The remaining panels display a selection of metal lines, overlaid with the Voigt profiles for the derived cloud model (in red) and blends (in light blue). The red tick marks above the normalised continuum indicate the locations of the absorption components. For all panels, the y-axis scale is residual intensity. The normalised quasar continuum and zero-level are shown by the blue dashed and green dotted lines respectively.
4.4 Individual Objects

Table 4.4: Ion column densities of the DLA in J0311−1722 at \( z_{\text{abs}} = 3.734 \)

<table>
<thead>
<tr>
<th>Ion</th>
<th>Transitions used</th>
<th>( \log N(X)/\text{cm}^{-2} )</th>
</tr>
</thead>
<tbody>
<tr>
<td>H I</td>
<td>1025, 1215</td>
<td>20.30 ± 0.06</td>
</tr>
<tr>
<td>C II</td>
<td>1036, 1334</td>
<td>14.02 ± 0.08</td>
</tr>
<tr>
<td>C II*</td>
<td>1335</td>
<td>13.55 ± 0.06</td>
</tr>
<tr>
<td>N I</td>
<td>1200.2</td>
<td>( \leq 13.07^a )</td>
</tr>
<tr>
<td>O I</td>
<td>1302</td>
<td>14.70 ± 0.08</td>
</tr>
<tr>
<td>Si II</td>
<td>1260, 1304</td>
<td>13.31 ± 0.07</td>
</tr>
<tr>
<td>Fe II</td>
<td>1125</td>
<td>( \leq 13.76^a )</td>
</tr>
</tbody>
</table>

\( ^a \) 3\( \sigma \) limiting rest frame equivalent width.

In this table, I also provide upper limits for the column densities of several key ions that are undetected at the signal-to-noise of the data. Specifically, I calculate the 3\( \sigma \) limiting rest-frame equivalent width, \( W_0 \), over the velocity interval of absorption exhibited by the weakest transition, which in this case is Si II \( \lambda 1304 \). A 3\( \sigma \) upper limit to the undetected feature is then derived using the optically thin limit approximation, \( N = 1.13 \times 10^{20} \cdot W_0/\lambda^2 f \text{ cm}^{-2} \). For N I I use the undetected \( \lambda 1200.2 \) line to derive \( W_0(\text{N I}) \leq 13 \text{ mÅ} \), which implies \( \log N(\text{N I})/\text{cm}^{-2} \leq 13.07 \). Similarly for Fe II \( \lambda 1125 \), \( W_0(\text{Fe II}) \leq 10 \text{ mÅ} \) implies \( \log N(\text{Fe II})/\text{cm}^{-2} \leq 13.76 \).

4.4.2 J0831+3358: DLA at \( z_{\text{abs}} = 2.3036 \)

J0831+3358 was observed with HIRES on 2009 December 9 under good conditions with sub-arcsecond seeing. A 1.148" wide slit was used which, as measured in Chapter 5, delivered a spectral resolution of 7.3 km s\(^{-1}\) FWHM. Unfortunately, the Ly\( \alpha \) line at \( \lambda_{\text{obs}} \approx 4015 \text{ Å} \) falls in a gap between two of the CCDs on the HIRES detector mosaic. Thus, for this DLA, I adopt the H I column density \( \log [N(\text{H I})/\text{cm}^{-2}] = 20.25 \pm 0.15 \) derived by Penprase et al. (2010) from their observations of this QSO at \( R \approx 5000 \), which is sufficient to resolve the broad damped profile of the Ly\( \alpha \) line.

The metal-lines in these data are well fit by a model with two components of roughly equal strength separated by 14 km s\(^{-1}\) (see Figure 4.3). Details of the derived cloud model are presented in Table 4.2, with the associated column densities given in Table 4.5. The C II lines are saturated in this DLA, however, I have a clean measurement of \( N(\text{O I}) \) from a number of unsaturated O I lines. In Table 4.5 I also provide 3\( \sigma \) upper limits to the N I and S II column densities which are undetected at the signal-to-noise of these data.

4.4.3 J1001+0343: DLA at \( z_{\text{abs}} = 3.0784 \)

This QSO was observed with UVES in service mode on the nights of 2009 April 19 & 29, 2010 January 11 & 27 and 2010 February 7. The total integration time was 33700 s, yielding a S/N per pixel of \( \sim 40 \) at 5000 Å. The DLA inline to J1001+0343 was also investigated by Penprase et al.
Figure 4.3: Same as Figure 4.2 for a selection of metal lines associated with the \( z_{\text{abs}} = 2.30364 \) DLA towards J0831+3358.

Table 4.5: \textsc{Ion column densities of the DLA in J0831+3358 at} \( z_{\text{abs}} = 2.30364 \)

<table>
<thead>
<tr>
<th>Ion</th>
<th>Transitions used</th>
<th>( \log N(X)/\text{cm}^{-2} )</th>
</tr>
</thead>
<tbody>
<tr>
<td>H I</td>
<td>1215</td>
<td>( 20.25 \pm 0.15 )^a</td>
</tr>
<tr>
<td>N I</td>
<td>1199.5</td>
<td>( \leq 12.78 )^b</td>
</tr>
<tr>
<td>O I</td>
<td>988.5, 988.6, 988.7, 1039, 1302</td>
<td>14.93 ( \pm 0.05 )</td>
</tr>
<tr>
<td>Al II</td>
<td>1670</td>
<td>12.19 ( \pm 0.06 )</td>
</tr>
<tr>
<td>Si II</td>
<td>989, 1193, 1260, 1304</td>
<td>13.75 ( \pm 0.04 )</td>
</tr>
<tr>
<td>S II</td>
<td>1250</td>
<td>( \leq 13.75 )^b</td>
</tr>
<tr>
<td>Fe II</td>
<td>1608</td>
<td>13.33 ( \pm 0.06 )</td>
</tr>
</tbody>
</table>

\(^a\)Penprase et al. (2010)

\(^b\)3\(\sigma\) limiting rest frame equivalent width.
Figure 4.4: Same as Figure 4.2, for the DLA towards J1001+0343 (black histogram) which exhibits a damped Ly$\alpha$ line at $z_{\text{abs}} = 3.07841$ (top panel). Here, the red continuous line shows the theoretical Voigt profile for an H I column density $\log[N(\text{H I})/\text{cm}^{-2}] = 20.21$. The remaining panels display a selection of metal lines. Note the different y-axis scale that is used for the weak Fe II $\lambda$1608 line (bottom-right panel).
Table 4.6: Ion column densities of the DLA in J1001+0343 at $z_{\text{abs}} = 3.07841$

<table>
<thead>
<tr>
<th>Ion</th>
<th>Transitions used</th>
<th>$\log N(X) / \text{cm}^{-2}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>H I</td>
<td>1215</td>
<td>$20.21 \pm 0.05$</td>
</tr>
<tr>
<td>C II</td>
<td>1036, 1334</td>
<td>$13.58 \pm 0.02$</td>
</tr>
<tr>
<td>N I</td>
<td>1200.2</td>
<td>$\leq 12.50^a$</td>
</tr>
<tr>
<td>O I</td>
<td>1039, 1302</td>
<td>$14.25 \pm 0.02$</td>
</tr>
<tr>
<td>Si II</td>
<td>1190, 1193, 1260, 1304, 1526</td>
<td>$12.86 \pm 0.01$</td>
</tr>
<tr>
<td>S II</td>
<td>1253</td>
<td>$\leq 12.91^a$</td>
</tr>
<tr>
<td>Fe II</td>
<td>1608</td>
<td>$12.50 \pm 0.14$</td>
</tr>
</tbody>
</table>

$^a$3σ limiting rest frame equivalent width.

(2010), being amongst the most metal-poor in their sample. From these observations I derive an H I column density of $\log [N(\text{H I}) / \text{cm}^{-2}] = 20.21 \pm 0.05$ from the wings of the Ly$\alpha$ absorption line. This compares well with the estimate by Penprase et al. (2010), $\log [N(\text{H I}) / \text{cm}^{-2}] = 20.15 \pm 0.10$. I present the Voigt profile fit to the Ly$\alpha$ line in the top panel of Figure 4.4.

The remaining panels of Figure 4.4 showcase a number of the available absorption lines that were used in deriving the cloud model. In fact, all of the available metal absorption lines are unsaturated in this DLA, thus providing reliable measurements of the elemental abundances. Several metal-line transitions of varying strength are well fit by a single component cloud model with a Doppler parameter of $b = 7.0 \pm 0.1 \text{ km s}^{-1}$ (see Table 4.2). The curve of growth analysis by Penprase et al. (2010) yielded a Doppler parameter of 7.5 km s$^{-1}$, which is in good agreement with that found here. The derived column densities for all available ions are presented in Table 4.6. I also provide 3σ upper limits to the N I and S II column densities which are undetected at the S/N of these data. Fe II $\lambda 1608$ is detected at the 3.6σ level. The corresponding fit is presented in the bottom-right panel of Figure 4.4 (note the different y-axis scale).

4.4.4 J1037+0139: DLA at $z_{\text{abs}} = 2.70487$

This QSO was also observed in service mode with UVES on 2010 February 12–14 and again on 2010 March 5. J1037+0139 was also one of the QSOs observed independently by Penprase et al. (2010). It is one of the faintest QSOs in this sample, requiring a total of 26075 s of integration to achieve a S/N $\sim 40$ at 5000 Å. The Ly$\alpha$ line falls on the edge of the blue detector, but this does not affect the accuracy of the H I column density since the blue wing of the damped Ly$\alpha$ line is still intact, and can be fit using the redshift derived from the well-defined narrow metal absorption lines. Moreover, I have access to Ly$\beta$, which also exhibits damping wings (although not as strong as that of Ly$\alpha$). I derive an H I column density of $\log [N(\text{H I}) / \text{cm}^{-2}] = 20.50 \pm 0.08$, which is consistent with estimates derived from the SDSS spectrum ($\log [N(\text{H I}) / \text{cm}^{-2}] = 20.45$; Prochaska & Wolfe 2009), as well as that derived by Penprase et al. (2010) ($\log [N(\text{H I}) / \text{cm}^{-2}] = 20.40 \pm 0.25$). I present the Voigt profile fit to the Ly$\alpha$ line in the top panel of Figure 4.5.

Again, the metal absorption is concentrated in a single component, in this case with a
4.4 Individual Objects

Figure 4.5: Same as Figure 4.2, for the DLA towards J1037+0139 (black histogram) which exhibits a damped Lyα line at $z_{\text{abs}} = 2.70487$ (top panel). The red continuous line shows the theoretical Voigt profile for an H\textsc{i} column density $\log[N(\text{H}\textsc{i})/\text{cm}^{-2}] = 20.50$. The remaining panels display a selection of metal lines. Note that the feature on the red wing of the strong Si\textsc{ii} 1260 absorption feature is due to Fe\textsc{ii} 1260 absorption.
Doppler parameter of 5.9 km s\(^{-1}\). The corresponding column densities for all of the available ions are listed in Table 4.7. For this system, both C\(\text{II}\)\(\lambda\)1036 and \(\lambda\)1334 are saturated and blended; however, a robust measure of the O\(\text{I}\) and Fe\(\text{II}\) column densities is afforded from several unsaturated transitions. I also have a clear detection of the N\(\text{I}\) triplet near \(\lambda_0 = 1200\,\text{Å}\).

### Table 4.7: Ion column densities of the DLA in J1037+0139 at \(z_{\text{abs}} = 2.70487\)

<table>
<thead>
<tr>
<th>Ion</th>
<th>Transitions used</th>
<th>(\log N(X)/\text{cm}^{-2})</th>
</tr>
</thead>
<tbody>
<tr>
<td>H(\text{I})</td>
<td>1025, 1215</td>
<td>20.50 ± 0.08</td>
</tr>
<tr>
<td>N(\text{I})</td>
<td>1199.5, 1200.2, 1200.7</td>
<td>13.27 ± 0.04</td>
</tr>
<tr>
<td>O(\text{I})</td>
<td>1039, 1302</td>
<td>15.06 ± 0.04</td>
</tr>
<tr>
<td>Al(\text{II})</td>
<td>1670</td>
<td>12.32 ± 0.03</td>
</tr>
<tr>
<td>Si(\text{II})</td>
<td>1260, 1304</td>
<td>13.97 ± 0.03</td>
</tr>
<tr>
<td>Fe(\text{II})</td>
<td>1143, 1144, 1608</td>
<td>13.53 ± 0.02</td>
</tr>
</tbody>
</table>

4.4.5 J1340+1106: DLA at \(z_{\text{abs}} = 2.50792\)

This QSO has previously been observed with both UVES (at a spectral resolution of 10.3 km s\(^{-1}\) FWHM; Ledoux, Petitjean, & Srianand 2003) and HIRES (at a spectral resolution of 8.1 km s\(^{-1}\) FWHM; Prochaska et al. 2003). However, given that this QSO intersects two VMP DLAs, one of which had potentially unsaturated C\(\text{II}\) lines, I decided to reobserve it with UVES for 29800 s at a slightly higher spectral resolution of 7.3 km s\(^{-1}\), and obtained complete spectral coverage from 3500 Å to almost 1μm.

Since the broad damped profile of the Ly\(\alpha\) line is independent of the spectral resolution, I combined the three datasets to obtain a high S/N ratio near the damped Ly\(\alpha\) line, from which I derived \(\log[N(\text{H}\text{I})/\text{cm}^{-2}] = 20.09 ± 0.05\). This model fit, along with the combined data, is shown in the top panel of Figure 4.6. The profiles of the metal absorption lines, on the other hand, are narrow (FWHM \(\lesssim 10\) km s\(^{-1}\)); therefore, their observed profiles are not independent of the spectral resolution. Thus, I separately combined the data of equal spectral resolution and individually read these three reduced spectra into \textsc{vpfit}, which convolved the fitted model with the spectral resolution appropriate to the data. The upshot of proceeding in this way is that the cloud model is then largely driven by the dataset of highest S/N for each absorption line that is input. However, since I cannot combine all three datasets, in Figure 4.6 I only present the dataset with the highest S/N near each absorption line.

Unfortunately, both of the C\(\text{II}\) absorption lines are saturated, and partially arise from nearby ionised gas, exhibiting a similar profile shape to the Al\(\text{II}\)\(\lambda\)1670 line. In fact, most of the metal absorption lines exhibit a second component on the blue wing (at \(v = -23\) km s\(^{-1}\) relative to the main component at \(z_{\text{abs}} = 2.507921\)), which appears to arise from ionised gas. This is confirmed by the absence of this blue component in O\(\text{I}\) absorption (see O\(\text{I}\)\(\lambda\)1302, and note that the absorption on the blue wing of the O\(\text{I}\)\(\lambda\)1039 line is due to unrelated absorption), which
Figure 4.6: Same as Figure 4.2, for the DLA towards J1340+1106 (black histogram) which exhibits a damped Lyα line at $z_{\text{abs}} = 2.50792$ (top panel). The red continuous line shows the theoretical Voigt profile for an H I column density $\log [N(\text{H} I)/\text{cm}^{-2}] = 20.09$. The remaining panels display a selection of metal lines.
Chapter 4. A survey for the most metal-poor DLAs

Table 4.8: Ion column densities of the DLA in J1340+1106 at $z_{\text{abs}} = 2.50792$

<table>
<thead>
<tr>
<th>Ion</th>
<th>Transitions used</th>
<th>$\log N(X)/\text{cm}^{-2}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>H I</td>
<td>1215</td>
<td>$20.09 \pm 0.05$</td>
</tr>
<tr>
<td>N I</td>
<td>1199.5, 1200.2</td>
<td>$12.83 \pm 0.05$</td>
</tr>
<tr>
<td>O I</td>
<td>1039, 1302</td>
<td>$15.05 \pm 0.03$</td>
</tr>
<tr>
<td>Al II</td>
<td>1670</td>
<td>$12.29 \pm 0.02$</td>
</tr>
<tr>
<td>Al III</td>
<td>1854, 1862</td>
<td>$11.36 \pm 0.15$</td>
</tr>
<tr>
<td>Si II</td>
<td>1190, 1193, 1304, 1526</td>
<td>$13.76 \pm 0.02$</td>
</tr>
<tr>
<td>Fe II</td>
<td>1096, 1143, 2344, 2374, 2382</td>
<td>$13.51 \pm 0.02$</td>
</tr>
</tbody>
</table>

has long been known to accurately trace neutral gas (Field & Steigman, 1971). The presence of ionised gas is also confirmed by the higher Al III/Al II ratio exhibited by the blue component.

I derived the cloud model for both of these components from the host of available Si II and Fe II lines, with additional constraints coming from the two O I lines. In Table 4.2 I present the fitting results from both the blue component (component 1; which arises from nearby mildly ionised gas), and the main component (component 2) which I attribute to the DLA. Fixing the parameters of this cloud model, I then derived the column density for all ions in both components. In Table 4.8, however, I only provide the column density for the single, dominant component that I attribute to the DLA. The model fits to the data are presented in the lower panels of Figure 4.6 where, as stated above, I only present the model and data that correspond to the highest S/N for each absorption line.

The detection of ions that arise in ionised gas, such as Al III, will later provide a useful means to test the accuracy of one of my underlying assumptions; that I can use the single dominant ion for each element to measure the elemental abundances in DLAs (see Section 4.5.1).

4.4.6 J1340+1106: DLA at $z_{\text{abs}} = 2.79583$

I now report on the second VMP DLA that is intersected by this QSO. Again, I treat the Ly$\alpha$ line and the metal lines of this DLA as detailed in Section 4.4.5. From the combined data I derive an H I column density of $\log[N(\text{H I})/\text{cm}^{-2}] = 21.00 \pm 0.06$. The combined spectrum in the region of the Ly$\alpha$ line, along with the profile fit, is reproduced in the top panel of Figure 4.7.

Turning now to the metal absorption lines, I have found that a two component cloud model (with Doppler parameters of $9.3 \pm 0.2$ and $6.45 \pm 0.05$ km s$^{-1}$ separated by $22$ km s$^{-1}$) provides a good fit to the data. This cloud model is perhaps the best determined in my dataset, given the high S/N of the data, and the numerous atomic transitions available. The Voigt profile fits to the metal lines are shown in the lower panels of Figure 4.7. However, as discussed in Section 4.4.5, I only present the model and data that correspond to the highest S/N for each absorption line. The column densities for all available ions are provided in Table 4.9. In this table, I have also provided the column densities for Al III and N II, which are coincident with the DLA, but are typically associated with H II regions.
Figure 4.7: Same as Figure 4.2, for the DLA towards J1340+1106 (black histogram) which exhibits a damped Ly$\alpha$ line at $z_{\text{abs}} = 2.79583$ (top panel). The red continuous line shows the theoretical Voigt profile for an H I column density $\log[N(\text{H} I)/\text{cm}^{-2}] = 21.00$. The remaining panels display a selection of metal lines. Continued next page.
Chapter 4. A survey for the most metal-poor DLAs

Velocity Relative to $z_{abs} = 2.7958285$ (km s$^{-1}$)
Table 4.9: Ion column densities of the DLA in J1340+1106 at $z_{\text{abs}} = 2.79583$

<table>
<thead>
<tr>
<th>Ion</th>
<th>Transitions used</th>
<th>$\log N(X)/\text{cm}^{-2}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>H I</td>
<td>1215</td>
<td>21.00 ± 0.06</td>
</tr>
<tr>
<td>C II*</td>
<td>1335</td>
<td>12.82 ± 0.07</td>
</tr>
<tr>
<td>N I</td>
<td>1134.4, 1134.9, 1199.5, 1200.7</td>
<td>14.08 ± 0.02</td>
</tr>
<tr>
<td>N II</td>
<td>1083</td>
<td>12.80 ± 0.10</td>
</tr>
<tr>
<td>O I</td>
<td>925, 976, 1302</td>
<td>16.04 ± 0.04</td>
</tr>
<tr>
<td>Al II</td>
<td>1670</td>
<td>13.27 ± 0.03</td>
</tr>
<tr>
<td>Al III</td>
<td>1854, 1862</td>
<td>12.19 ± 0.07</td>
</tr>
<tr>
<td>Si II</td>
<td>1020, 1193, 1260, 1304, 1526, 1808</td>
<td>14.70 ± 0.02</td>
</tr>
<tr>
<td>S II</td>
<td>1253, 1259</td>
<td>14.30 ± 0.02</td>
</tr>
<tr>
<td>Ar I</td>
<td>1048, 1066</td>
<td>13.14 ± 0.02</td>
</tr>
<tr>
<td>Cr II</td>
<td>2056, 2062</td>
<td>12.61 ± 0.01</td>
</tr>
<tr>
<td>Fe II</td>
<td>1063.1, 1081, 1096, 1125, 1608, 1611, 2344, 2382</td>
<td>14.32 ± 0.01</td>
</tr>
<tr>
<td>Ni II</td>
<td>1317, 1370, 1454, 1709, 1741, 1751</td>
<td>13.09 ± 0.03</td>
</tr>
</tbody>
</table>

Table 4.10: Ion column densities of the DLA in J1419+0829 at $z_{\text{abs}} = 3.04973$

<table>
<thead>
<tr>
<th>Ion</th>
<th>Transitions used</th>
<th>$\log N(X)/\text{cm}^{-2}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>H I</td>
<td>1025, 1215</td>
<td>20.40 ± 0.03</td>
</tr>
<tr>
<td>N I</td>
<td>1199.5, 1200.2, 1200.7</td>
<td>13.28 ± 0.02</td>
</tr>
<tr>
<td>O I</td>
<td>936, 948, 976, 988.5, 988.6, 988.7, 1039, 1302</td>
<td>15.17 ± 0.02</td>
</tr>
<tr>
<td>Si II</td>
<td>989, 1190, 1193, 1260, 1304, 1526</td>
<td>13.83 ± 0.01</td>
</tr>
<tr>
<td>Fe II</td>
<td>1063.1, 1608</td>
<td>13.54 ± 0.03</td>
</tr>
</tbody>
</table>

4.4.7 J1419+0829: DLA at $z_{\text{abs}} = 3.04973$

The spectrum of J1419+0829 ($z_{\text{em}} = 3.034$) was recorded for 29800 s with UVES in service mode, resulting in a S/N near 5000 Å (near the red wing of the damped Ly$\alpha$ line) of $\sim 40$. This high S/N in combination with a virtually uninterrupted red wing to the Ly$\alpha$ absorption allows a very accurate measurement of the H I column density, $\log[N(\text{H I})/\text{cm}^{-2}] = 20.40 ± 0.03$. The Voigt profile fit to the Ly$\alpha$ line is shown in the top panel of Figure 4.8, with a selection of the associated metal absorption lines in the remaining panels.

A cloud model with two components separated by $\sim 14$ km s$^{-1}$ (Table 4.2) provides a good fit to the data. This cloud model is well determined by a host of O I and Si II lines (see Table 4.10). Both C II $\lambda 1334$ and $\lambda 1036$ are saturated; column densities for N I, O I, Si II, and Fe II are listed in Table 4.10.

4.4.8 The final VMP DLA sample

In order to augment this survey for VMP DLAs, I have searched the literature for known examples satisfying the following conditions: (i) the QSO spectra were observed at high spectral resolution ($R > 30000$) – in practice this meant that the data were recorded with either UVES or HIRES; (ii) $[\text{Fe/H}] \leq -2.0$; and (iii) at least one unsaturated O I absorption line from which
Figure 4.8: Same as Figure 4.2, for the DLA towards J1419+0829 (black histogram) which exhibits a damped Ly$\alpha$ line at $z_{\text{abs}} = 3.04973$ (top left panel). The red continuous line shows the theoretical Voigt profile for an H I column density $\log [N(\text{H} I)/\text{cm}^{-2}] = 20.40$. The remaining panels display a selection of metal lines.
[O/H] could be measured. These conditions were imposed to select a sample of measurements from the literature which is highly compatible to the data presented here and whose metal abundances could be adopted without reanalysing the spectra (although I referred all measurements to the same solar abundance scale – see Appendix A). The literature trawl yielded an additional ten DLAs satisfying the above conditions; together with the observations reported here and in Pettini et al. (2008a), they form a sample of 22 VMP DLAs.

Relevant details of the full sample are collected in Table 4.11, where I list absorption redshifts, neutral hydrogen column densities and element abundances for a selection of the metals most commonly observed in VMP DLAs, including: (1) C, N, O – the first elements synthesised in the chain of stellar nucleosynthesis; (2) Al – an odd atomic number element; (3) Si – an even atomic number element; and (4) Fe – an iron-peak element. The uncertainty in each abundance includes the error in H\textsc{i}. In Appendix B, I provide a similar table listing the column densities of each ion from which these abundances were derived.

4.5 Abundance Analysis

As in previous DLA work, I assume that each element resides in a single dominant ionisation stage in the neutral gas. Thus, the abundance of a given element is found by taking the ratio of the dominant ions column density to that of H\textsc{i}, and referring it to a solar scale (i.e. \([X/H] = \log(N(X)/N(H\textsc{i})) - \log(X/H)_{\odot}\); see Appendix A). Thus, measuring elemental abundances in VMP DLAs is a relatively straightforward process. Some uncertainty arises, however, if the observed metal-line absorption from ions that are dominant in H\textsc{i} regions, does not perfectly trace the H\textsc{i} gas. For example, if the dominant ion for a given element is also present in nearby H\textsc{ii} gas (often the case for singly ionised species; e.g. C\textsc{ii}, Si\textsc{ii}, Fe\textsc{ii}), I will overestimate the abundance of this element. On the other hand, if the dominant ion for a given element is mildly ionised in the H\textsc{i} gas itself (usually the case for neutral species; e.g. N\textsc{i}), the element’s abundance will be under-estimated. In addition to these ionisation corrections, further uncertainties may be introduced into the abundance analysis if some refractory elements have condensed to form dust grains. Whilst both of these concerns are expected to be negligible in the VMP regime (Vladilo et al., 2001; Vladilo, 2004), I reassess their importance in the following sub-sections.

4.5.1 Ionisation Corrections

Ionisation corrections are known to be small when the neutral hydrogen column density is in excess of \(~10^{20}\) atoms cm\(^{-2}\). Nevertheless, for each element X with the dominant ionisation stage N, one needs to apply a small correction, IC(X), to recover the true elemental abundance,

\[ [X/H] = [X_N/H\textsc{i}] + IC(X). \]  (4.1)
### Table 4.11: C, N, O, Al, Si, and Fe Abundance Measurements in VMP DLAs

<table>
<thead>
<tr>
<th>QSO</th>
<th>$z_{\text{abs}}$</th>
<th>$\log N(\text{H})$ (cm$^{-2}$)</th>
<th>[C/H]</th>
<th>[N/H]</th>
<th>[O/H]</th>
<th>[Al/H]</th>
<th>[Si/H]</th>
<th>[Fe/H]</th>
<th>Ref.$^a$</th>
</tr>
</thead>
<tbody>
<tr>
<td><strong>The VMP DLA Sample</strong></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>J0035$-$0918</td>
<td>2.34010</td>
<td>20.55 ± 0.10</td>
<td>−1.51 ± 0.18</td>
<td>−2.33 ± 0.12</td>
<td>−2.28 ± 0.13</td>
<td>−3.26 ± 0.11</td>
<td>−2.65 ± 0.11</td>
<td>−3.04 ± 0.12</td>
<td>2</td>
</tr>
<tr>
<td>J0311$-$1722</td>
<td>3.73400</td>
<td>20.30 ± 0.06</td>
<td>−2.71 ± 0.10</td>
<td>≤ −2.52</td>
<td>−2.29 ± 0.10</td>
<td>...</td>
<td>−2.50 ± 0.09</td>
<td>≤ −2.01</td>
<td>1</td>
</tr>
<tr>
<td>J0831$+$3358</td>
<td>2.30364</td>
<td>20.25 ± 0.15</td>
<td>...</td>
<td>≤ −2.76</td>
<td>−2.01 ± 0.16</td>
<td>−2.50 ± 0.16</td>
<td>−2.01 ± 0.16</td>
<td>−2.39 ± 0.16</td>
<td>1,4</td>
</tr>
<tr>
<td>Q0913$+$072</td>
<td>2.61843</td>
<td>20.34 ± 0.04</td>
<td>−2.79 ± 0.06</td>
<td>−3.34 ± 0.13</td>
<td>−2.40 ± 0.04</td>
<td>−3.00 ± 0.05</td>
<td>−2.55 ± 0.04</td>
<td>−2.82 ± 0.04</td>
<td>3</td>
</tr>
<tr>
<td>J1001$+$0343</td>
<td>3.07841</td>
<td>20.21 ± 0.05</td>
<td>−3.06 ± 0.05</td>
<td>≤ −3.00</td>
<td>−2.65 ± 0.05</td>
<td>...</td>
<td>−2.86 ± 0.05</td>
<td>−3.18 ± 0.15</td>
<td>1</td>
</tr>
<tr>
<td>J1016$+$4040</td>
<td>2.81633</td>
<td>19.90 ± 0.11</td>
<td>−2.67 ± 0.12</td>
<td>≤ −2.43</td>
<td>−2.46 ± 0.11</td>
<td>...</td>
<td>−2.51 ± 0.12</td>
<td>...</td>
<td>3</td>
</tr>
<tr>
<td>J1037$+$0139</td>
<td>2.70487</td>
<td>20.50 ± 0.08</td>
<td>...</td>
<td>−2.52 ± 0.09</td>
<td>−2.13 ± 0.09</td>
<td>−2.62 ± 0.09</td>
<td>−2.04 ± 0.09</td>
<td>−2.44 ± 0.08</td>
<td>1</td>
</tr>
<tr>
<td>J1340$+$1106</td>
<td>2.50792</td>
<td>20.09 ± 0.05</td>
<td>...</td>
<td>−2.55 ± 0.07</td>
<td>−1.73 ± 0.06</td>
<td>−2.24 ± 0.05</td>
<td>−1.84 ± 0.05</td>
<td>−2.05 ± 0.05</td>
<td>1</td>
</tr>
<tr>
<td>J1340$+$1106</td>
<td>2.79583</td>
<td>21.00 ± 0.06</td>
<td>...</td>
<td>−2.21 ± 0.06</td>
<td>−1.65 ± 0.07</td>
<td>−2.17 ± 0.07</td>
<td>−1.81 ± 0.06</td>
<td>−2.15 ± 0.06</td>
<td>1</td>
</tr>
<tr>
<td>J1419$+$0829</td>
<td>3.04973</td>
<td>20.43 ± 0.03</td>
<td>...</td>
<td>−2.41 ± 0.04</td>
<td>−1.92 ± 0.04</td>
<td>...</td>
<td>−2.08 ± 0.03</td>
<td>−2.33 ± 0.04</td>
<td>1</td>
</tr>
<tr>
<td>J1558$+$4053</td>
<td>2.55322</td>
<td>20.30 ± 0.04</td>
<td>−2.51 ± 0.07</td>
<td>−2.93 ± 0.08</td>
<td>−2.45 ± 0.06</td>
<td>−2.82 ± 0.07</td>
<td>−2.49 ± 0.04</td>
<td>−2.70 ± 0.07</td>
<td>3</td>
</tr>
<tr>
<td>Q2206$-$199</td>
<td>2.07624</td>
<td>20.43 ± 0.04</td>
<td>−2.45 ± 0.05</td>
<td>−2.93 ± 0.06</td>
<td>−2.07 ± 0.05</td>
<td>−2.69 ± 0.04</td>
<td>−2.29 ± 0.04</td>
<td>−2.57 ± 0.04</td>
<td>3</td>
</tr>
<tr>
<td><strong>Literature VMP DLAs</strong></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>Q0000$-$2620</td>
<td>3.39012</td>
<td>21.41 ± 0.08</td>
<td>...</td>
<td>−2.00 ± 0.08</td>
<td>−1.68 ± 0.13</td>
<td>...</td>
<td>−1.86 ± 0.08</td>
<td>−2.01 ± 0.09</td>
<td>5</td>
</tr>
<tr>
<td>Q0112$-$306</td>
<td>2.41844</td>
<td>20.50 ± 0.08</td>
<td>...</td>
<td>−2.63 ± 0.09</td>
<td>−2.24 ± 0.11</td>
<td>...</td>
<td>−2.39 ± 0.08</td>
<td>−2.64 ± 0.09</td>
<td>6</td>
</tr>
<tr>
<td>J0140$-$0839</td>
<td>3.69660</td>
<td>20.75 ± 0.15</td>
<td>−3.05 ± 0.17</td>
<td>≤ −3.66</td>
<td>−2.75 ± 0.15</td>
<td>−3.37 ± 0.16</td>
<td>−2.75 ± 0.17</td>
<td>−3.45 ± 0.24$^b$</td>
<td>7</td>
</tr>
<tr>
<td>J0307$-$4945</td>
<td>4.46658</td>
<td>20.67 ± 0.09</td>
<td>...</td>
<td>−2.39 ± 0.15</td>
<td>−1.45 ± 0.19</td>
<td>−1.75 ± 0.11</td>
<td>−1.50 ± 0.11</td>
<td>−1.93 ± 0.19</td>
<td>8</td>
</tr>
<tr>
<td>Q1108$-$077</td>
<td>3.60767</td>
<td>20.37 ± 0.07</td>
<td>...</td>
<td>≤ −2.82</td>
<td>−1.69 ± 0.08</td>
<td>...</td>
<td>−1.54 ± 0.07</td>
<td>−1.96 ± 0.07</td>
<td>6</td>
</tr>
<tr>
<td>J1337$+$3153</td>
<td>3.16768</td>
<td>20.41 ± 0.15</td>
<td>−2.86 ± 0.16</td>
<td>≤ −2.90</td>
<td>−2.67 ± 0.17</td>
<td>−2.85 ± 0.16</td>
<td>−2.68 ± 0.16</td>
<td>−2.74 ± 0.30</td>
<td>9</td>
</tr>
<tr>
<td>J1558$-$0031</td>
<td>2.70262</td>
<td>20.67 ± 0.05</td>
<td>...</td>
<td>−1.50 ± 0.05$^c$</td>
<td>−1.50 ± 0.05$^c$</td>
<td>...</td>
<td>−1.94 ± 0.05$^c$</td>
<td>−2.03 ± 0.05$^c$</td>
<td>10</td>
</tr>
<tr>
<td>Q1946$+$7658</td>
<td>2.84430</td>
<td>20.27 ± 0.06</td>
<td>...</td>
<td>−2.97 ± 0.07</td>
<td>−2.14 ± 0.06</td>
<td>...</td>
<td>−2.18 ± 0.06</td>
<td>−2.50 ± 0.06</td>
<td>11</td>
</tr>
<tr>
<td>Q2059$-$360</td>
<td>3.08293</td>
<td>20.98 ± 0.08</td>
<td>...</td>
<td>−2.32 ± 0.08</td>
<td>−1.58 ± 0.09</td>
<td>...</td>
<td>−1.63 ± 0.09</td>
<td>−1.97 ± 0.08</td>
<td>6</td>
</tr>
<tr>
<td>J2155$+$1358</td>
<td>4.21244</td>
<td>19.61 ± 0.10</td>
<td>−2.09 ± 0.12</td>
<td>...</td>
<td>−1.80 ± 0.11</td>
<td>−2.13 ± 0.20</td>
<td>−1.87 ± 0.11</td>
<td>−2.15 ± 0.25</td>
<td>12</td>
</tr>
</tbody>
</table>

$^a$References—1: This work; 2: Chapter 5; 3: Pettini et al. (2008a); 4: Penprase et al. (2010); 5: Molaro et al. (2000); 6: Petitjean, Ledoux, & Srianand (2008); 7: Ellison et al. (2010); 8: Dessauges-Zavadsky et al. (2001); 9: Srianand et al. (2010); 10: O’Meara et al. (2006); 11: Prochaska et al. (2002); 12: Dessauges-Zavadsky et al. (2003). $^b$Ellison et al. (2010) quote a 3σ upper limit to the Fe II column density of log $N$(Fe II)/cm$^{-2}$ < 12.73. I have since rereduced these data (as described in Section 4.2), and detected the Fe II λ1608 line at the 4σ level. The Fe abundance quoted here is derived using the optically thin limit approximation to measure $N$(Fe II). $^c$The metal ion uncertainty for this measurement was not provided by the authors. Thus, I only quote the uncertainty in $N$(H I).
4.5 Abundance Analysis

Figure 4.9: Ionisation corrections for two typical VMP DLAs, respectively with low and high H I column density. **Left Panels:** The ionisation corrections, as defined by Equation 4.1, for the most commonly observed elements in VMP DLAs are plotted as a function of gas volume density. The vertical dashed lines correspond to the estimated gas densities of the two DLAs (shown by the solid vertical line in the right panels). **Right Panels:** Column density ratios for successive ion stages of Al (blue squares connected by a solid line) and N (red diamonds connected by a solid line) vs. gas density. I also plot the observed values of the Al \( \text{Al}^{\text{III}}/\text{Al}^{\text{II}} \) and \( \text{N}^{\text{II}}/\text{N}^{\text{I}} \) ratios (solid horizontal lines) along with their uncertainties (dotted lines) for the two VMP DLAs used as examples in this test (see text for further details).

To estimate the magnitude of such corrections, I used the CLOUDY photoionisation software developed by Ferland et al. (1998) to model two of the VMP DLAs from this sample, chosen to represent the range of \( N(\text{H}^\text{I}) \) values that is reported here: one DLA with a low \( N(\text{H}^\text{I}) \) column density (the DLA with \( \log[N(\text{H}^\text{I})/\text{cm}^{-2}] = 20.09 \) towards J1340+1106 at \( z_{\text{abs}} = 2.50792 \)) and the other with a high \( N(\text{H}^\text{I}) \) (the DLA with \( \log[N(\text{H}^\text{I})/\text{cm}^{-2}] = 21.00 \) towards J1340+1106 at \( z_{\text{abs}} = 2.79583 \)). In both cases, I modelled the VMP DLA as a plane-parallel slab of constant volume density gas in the range \( -3 < \log[n(\text{H})/\text{cm}^{-3}] < 3 \), irradiated by the cosmic microwave background and UV background at the appropriate redshift. Using the solar abundance scale in Appendix A I globally scaled the metal abundances of the VMP DLA to be \( 10^{-2} Z_\odot \). The simulations were stopped once the \( \text{H}^\text{I} \) column density of the DLA was reached, at which point the simulated ion column densities of the slab were output.

Once the value of the background radiation is assumed, the ionisation correction for each element depends on the volume density of the gas (left panels of Figure 4.9), which may be estimated by considering the ratio of successive ion stages (right panels of Figure 4.9). For the low \( N(\text{H}^\text{I}) \) DLA being considered as an example here, I measure \( N(\text{Al}^{\text{III}})/N(\text{Al}^{\text{II}}) = -0.93 \pm \)
0.15, implying a gas density of \( \log \left[ \frac{n(H)}{\text{cm}^{-3}} \right] \approx -0.7 \pm 0.4 \). For the high \( N(\text{H} I) \) DLA, I measure \( \frac{N(\text{Al} III)}{N(\text{Al} II)} = -1.08 \pm 0.08 \) and \( \frac{N(\text{N} II)}{N(\text{N} I)} = -1.28 \pm 0.10 \), which are both consistent with a gas density of \( \log \left[ \frac{n(H)}{\text{cm}^{-3}} \right] \approx -1.0 \pm 0.2 \). At these values of the gas density, it can be seen that the ionisation corrections for the VMP DLAs in this sample are \( \lesssim 0.1 \) dex for the main elements of interest. In the analysis that follows, I have therefore not corrected any of the measured abundances for ionisation effects.

### 4.5.2 Dust Depletion

To measure accurately element abundances in DLAs, the fraction of a given element that is not observed in the gas phase, but is instead locked up in dust grains, must also be considered. To account for this effect, one ideally considers the relative abundances of a refractory and a volatile element (e.g. \([\text{Cr/Zn}]\)), and compares this to the expected intrinsic nucleosynthetic ratio (typically the ratio seen in stars of comparable metallicity). Previous studies based on such a comparison have shown that DLAs exhibit minimal dust depletion when \([\text{Fe/H}] \lesssim -2\) (Pettini et al., 1997a; Akerman et al., 2005). Unfortunately, the \( \text{Cr} II \) and \( \text{Zn} II \) lines are too weak in the VMP regime to be measured, so one needs to resort to more abundant elements, such as \( \text{Si} \) and \( \text{Fe} \), which are known to be depleted to different degrees (Fe is more readily incorporated into dust grains than Si).

The most metal-poor stars in the halo of our Galaxy suggest that the intrinsic nucleosynthetic ratio of \( \text{Si}/\text{Fe} \) is virtually independent of metallicity, corresponding roughly to a constant of \([\text{Si}/\text{Fe}] = +0.37 \pm 0.15\) (Cayrel et al., 2004). Such a plateau was first seen in DLAs by Prochaska & Wolfe (2002), who found \([\text{Si}/\text{Fe}] \approx +0.3\) when \([\text{Fe/H}] < -2\) (see updated version in Wolfe et al. 2005). A similar study was also conducted by Vladilo (2002), who suggested there may still exist some mild depletion onto dust, resulting in a plateau of \([\text{Si}/\text{Fe}] \approx +0.25\). From the 19 VMP DLAs in this sample, I find that \([\langle \text{Si}/\text{Fe} \rangle] = +0.30 \pm 0.09\), which is certainly consistent with minimal dust depletion. I therefore proceed under the assumption that dust has a negligible effect on the derived elemental abundances.

### 4.6 Comparing VMP DLAs and stars

DLAs likely experience very different chemical histories to that of the stars in the halo of our Galaxy. However, in the limit of decreasing metallicity, both stars and DLAs are polluted by very few previous generations of stars. In fact, since the physical conditions of the gas that gives rise to DLAs are conducive to forming stars (Noterdaeme et al., 2008; Jorgenson et al., 2009), it certainly seems plausible that some of the VMP stars in the halo of our Galaxy originally condensed out of a VMP DLA. If this is indeed true, one would therefore expect both VMP stars and DLAs to share similar chemical signatures at the lowest metallicities. In the following subsections, I test the validity of this expectation, by directly comparing the abundances of the
observed in VMP DLAs in this survey to the most recent abundance measurements of VMP stars.

4.6 Comparing VMP DLAs and stars

4.6.1 Revisiting C/O at low metallicity

I first consider the trend of C/O at low metallicity, which has received a great deal of attention in recent years. The evolution of the C/O ratio when \([\text{O}/\text{H}] \gtrsim -1.0\) has been known for a while; \([\text{C}/\text{O}]\) increases linearly from \(\sim -0.6\) to solar with increasing \([\text{O}/\text{H}]\) (see Figure 4.10). This trend is thought to be due to the increased, metallicity-dependent, carbon yields of massive rotating stars, combined with the delayed release of carbon from low and intermediate mass stars (Akerman et al., 2004). Based on current models of Population II nucleosynthesis, below \([\text{O}/\text{H}] \sim -1.0\), \([\text{C}/\text{O}]\) is predicted to decrease (or perhaps plateau) with decreasing metallicity. Indeed, such a plateau was first reported for a sample of halo stars by Tomkin et al. (1992), who measured \([\text{C}/\text{O}]\) abundances down to \([\text{O}/\text{H}] \sim -1.7\).

This work was extended to even lower oxygen abundances by Akerman et al. (2004) and Spite et al. (2005). Contrary to the supposed decrease in \([\text{C}/\text{O}]\), these authors uncovered quite the opposite trend when \([\text{O}/\text{H}] \sim -2.0\); an extrapolation of this trend suggests that \([\text{C}/\text{O}]\) could reach near-solar values when \([\text{O}/\text{H}] \sim -3.0\). Three possibilities have been suggested to explain this behaviour: (1) the leftover signature of a high-carbon producing generation of Population III stars (Chieffi & Limongi, 2002; Umeda & Nomoto, 2003; Heger & Woosley, 2010); (2) a rapidly-rotating generation of low-metallicity Population II stars (Chiappini et al., 2006); or (3) systematic uncertainties in the adopted 1D LTE analysis. The third of these possibilities has recently been ruled out by Fabbian et al. (2009a), who conducted a non-LTE analysis of the lines used by Akerman et al. (2004), whilst including further constraints from additional carbon lines. Their analysis, however, is based on homogenous, plane-parallel (1D) model atmospheres rather than the more realistic inhomogenous (3D) models which have a somewhat different mean temperature structure. Nevertheless, it is unlikely that the use of 3D models will affect the derived \([\text{C}/\text{O}]\) significantly, since the abundances of C and O are both determined from high-excitation atomic lines, which react in the same way to changes in the temperature structure. The stellar C/O trend at low metallicity is therefore expected to be robust.

Such an ‘unexpected’ trend is perhaps not so surprising in hindsight, since several other lines of evidence support a high-carbon producing generation of early stars, including: (1) the fact that a high carbon abundance is required at early times to efficiently cool the gas, and drive the transition from Population III to Population II star formation (Frebel, Johnson, & Bromm, 2007); (2) observations of the three most iron-poor stars known to date have revealed that they all exhibit extreme carbon enhancements (Christlieb et al., 2002; Frebel et al., 2005; Norris et al., 2007); (3) in addition, the fraction of all carbon-enhanced metal-poor (CEMP) stars is thought to increase with decreasing metallicity (Beers & Christlieb, 2005); and (4) for at least a subset of these CEMP stars, it has been suggested that their carbon-enhancement reflects the
Chapter 4. A survey for the most metal-poor DLAs

Figure 4.10: C and O abundances in VMP DLAs (red triangles) observed at high spectral resolution. The green triangle represents a carbon-enhanced DLA that will be considered separately in Chapter 5. For comparison, I also plot the sample of metal-poor halo stars analysed by Fabbian et al. (2009a), where I adopt the values for C and O that are calculated assuming the Drawin (1969) formula for collisions with hydrogen ($S_H = 1$; blue squares). I also show a sample of thin- and thick-disc stars (filled and open circles respectively) from Bensby & Feltzing (2006). The red hatched region corresponds to the transition discriminant outlined by Frebel, Johnson, & Bromm (2007), where the uncertainty in this relation is shown by the light red shaded region. All of the plotted data have been referred to the Asplund et al. (2009) solar abundance scale.

The composition of the cloud of gas from which the CEMP star first condensed (Ryan et al., 2005; Aoki et al., 2007).

Additional evidence for a high carbon-producing early generation of stars has recently been provided by studies of the most metal-poor DLAs, which probe entire clouds of near-pristine gas. The first dedicated survey to uncover the chemical properties of these systems was conducted by Pettini et al. (2008a). These authors found that DLAs and stars tell the same story in the metal-poor regime – both show elevated, near-solar values of [C/O] when [O/H] $< -2.0$. In fact, the medium spectral resolution study by Penprase et al. (2010) suggests that [C/O] further increases to super-solar values at even lower metallicity. Further evidence for an increased carbon yield by an early generation of stars has recently come to light with the discovery of a metal-poor DLA that exhibits a C/Fe ratio 35 times greater than solar (this will be the subject of Chapter 5).

In Figure 4.10, I show the updated plot of [C/O] versus [O/H] for the full DLA survey, together with values for the metal-poor halo stars (blue squares) analysed by Fabbian et al. (2009a), and a sample of thin- and thick-disc stars with C and O abundances determined from forbidden lines (filled and open black circles respectively; Bensby & Feltzing 2006). All data shown in Figure 4.10 refer to cross sections for collisions with hydrogen atoms based on the classical recipe of Drawin (1969), i.e. using a scaling factor $S_H = 1$. If hydrogen collisions are completely neglected ($S_H = 0$), this trend is ‘stretched’ to higher [C/O] (by about 0.2 dex) and lower [O/H] (by about 0.3 dex at the lowest values of [O/H]).
data have been corrected for the updated Asplund et al. (2009) solar abundance scale (see Appendix A). I first note that there is generally a good agreement – both in the trend and the dispersion of [C/O] – between the most metal-poor stars and DLAs. The new DLA measurements reported here confirm the initial indications from the more limited samples considered by Pettini et al. (2008a) and Penprase et al. (2010). The main departure from this trend, represented by the green triangle in Figure 4.10, will be considered separately in Chapter 5. Aside from this system, no other DLA exhibits super-solar [C/O]. This statement is also true for the sample of seven O I absorbers at \( z_{\text{abs}} \sim 6 \) recently reported by Becker et al. (2011, see also Becker et al. 2006).

It is thus somewhat surprising that Penprase et al. (2010) found [C/O] \( \geq 0.0 \) for four out of the five VMP DLAs in which they could measure this ratio. The difference may be due to different sample criteria between this survey and theirs: whilst this survey has excluded absorption systems where the C II lines are saturated, such cases may be more difficult to recognise at the lower resolution of the ESI spectra analysed by Penprase et al. (2010). VMP DLAs with super-solar C/O ratios may well exist (and indeed the green triangle in Figure 4.10 is one such example), but their C abundance is generally more difficult to measure with confidence due to line saturation. On the other hand, this survey is not biased against uncovering systems with lower [C/O] values than those reported here; thus, the VMP DLAs in the current sample define a lower-envelope in the [C/O] versus [O/H] plane. This envelope appears to be in good agreement with the envelope defined by VMP stars (see Figure 4.10).

In Figure 4.10 I also show the ‘Frebel criterion’ (red hatched region; Frebel, Johnson, & Bromm 2007; see also Bromm & Loeb 2003) which states that, if the fine-structure lines of O I and C II dominate the cooling in a near-pristine cloud of gas that has been enriched to some critical metallicity, then the first low mass Population II stars will form. Thus, given these conditions, no Population II star should be observed in this red hatched region (where the uncertainty in this region is given by the light red shaded band). DLAs, on the other hand, are not restricted by this criterion. Indeed, if a cloud of gas was to be observed in this Population II star ‘forbidden zone’, it may very well form a collection of massive stars below or near the critical metallicity! Such systems, if found, would provide a unique window to study the transition from Population III to Population II star formation.

With these considerations in mind, I note that all DLAs in this sample will, perhaps unsurprisingly, form Population II stars. One might, therefore, be tempted to conclude that the stars in the halo of our Milky Way represent the same distribution that defines VMP DLAs. To test this possibility, I performed a linear fit to [C/O] versus [O/H] for the halo stars that have [O/H] \( \leq -2.0 \), and calculated the deviations about this line for both stars and DLAs. A Kolmogorov-Smirnov (K-S) test between the calculated deviations reveals a 75% chance that both VMP stars and DLAs are drawn from the same population, which is inconclusive given the present statistics. For this test I have used the C/O values in halo stars that were derived assuming efficient hydrogen collisions (with a scaling factor \( S_H = 1 \); see footnote 8). If instead
I adopt the stellar values derived for inefficient collisions \( S_H = 0 \), I find a 90% chance that both samples are drawn from the same population. Such good agreement between halo stars and DLAs does not support the recent claim by Tsujimoto & Bekki (2011), who suggest that the initial mass function (IMF) of the stars that enriched metal-poor DLAs is different to the IMF of the stars that enriched galactic halo stars with their metals. The good general agreement I report here between stars and DLAs – both exhibiting an elevated C/O ratio at the lowest metallicities probed – points to a universal origin for their C/O ‘excess’ in this regime.

**4.6.2 The O/Fe debate in the metal-poor regime**

I now turn to the relative abundances of oxygen and iron at low metallicity. Whilst I cannot do justice to the extensive literature on this topic, I outline below the basic facts that are relevant to this discussion, and direct the interested reader to the comprehensive review by McWilliam (1997).

The largest oxygen yield comes from the most massive stars that explode as SNe II. The budget for iron, on the other hand, is largely contributed by SNe Ia which typically explode \(~1\) Gyr later (see e.g. Greggio 2010).\(^9\) Therefore, at early times (when the metallicity is low), one expects O to be enhanced relative to Fe. At later times, when the delayed contribution of Fe from SNe Ia kicks in, there is a break in the O/Fe trend which is then expected to decrease. Thus, the relative abundance of O and Fe allows one to measure the relative contribution of SNe Ia and SNe II (see e.g. the qualitative discussion by Wheeler, Sneden, & Truran 1989).

In the Milky Way, the break in \([O/Fe]\) occurs roughly at \([\text{Fe/H}] \simeq -1\). Whilst there is sound agreement regarding the nature of the trend in \([O/Fe]\) when \([\text{Fe/H}] \gtrsim -1\), the behaviour of \([O/Fe]\) when \([\text{Fe/H}] \lesssim -1\) is less certain. This disagreement stems from the uncertainty of the oxygen abundances measured in metal-poor stars: there are four different indicators of the oxygen abundance, and to some extent they all disagree with one another in the metal-poor regime (García Pérez et al., 2006).

Perhaps the most reliable \([O/H]\) indicator at low metallicity is the forbidden \([O \text{ I}] \lambda 6300\) line which, despite being subject to 3D corrections of \(~-0.2\) dex when \([\text{Fe/H}] \sim -2\) (Nissen et al., 2002; Collet, Asplund, & Trampedach, 2007), is known to form in LTE (Asplund, 2005). Unfortunately, this line becomes very weak when \([\text{Fe/H}] < -2\) and its detection requires data of high S/N. After accounting for 3D corrections to the O abundance, most authors conclude that the O/Fe ratio is approximately constant at \([O/Fe] \simeq +0.4\) for \([\text{Fe/H}] \lesssim -1\) (Nissen et al., 2002; Cayrel et al., 2004; García Pérez et al., 2006), with perhaps a slight increase towards the lowest metallicities.

The most commonly used diagnostic for measuring \([O/H]\) in stars is the \([O \text{ I}]\) triplet near 777 nm \((\lambda = 7771.9, 7774.2, 7775.4 \text{ Å})\), despite the fact that it suffers from large non-LTE

---

\(^9\)Iron may also be contributed by SNe Ia that ‘promptly’ explode at early times \((\sim 0.1\) Gyr). For the relevant details, I direct the interested reader to the discussion by Mannucci, Della Valle, & Panagia (2006).
4.6 Comparing VMP DLAs and stars

corrections (Fabbian et al., 2009b). However, contrary to the nearly constant value of [O/Fe] below [Fe/H] \( \lesssim -1 \) deduced from the \([\text{O} \text{I}] \lambda 6300\) line, an LTE analysis of the O I triplet leads to a quasi-linear increase in [O/Fe] with decreasing metallicity (see e.g. Fulbright & Johnson 2003). This discrepancy is often blamed on the uncertain (negative) non-LTE corrections to the O I triplet. In order to make headway with the [O/Fe] conflict, Fabbian et al. (2009b) performed a detailed non-LTE analysis of the O I triplet, and found corrections amounting to \( \gtrsim 0.5\) dex when [Fe/H] = -3.0, increasing rapidly at lower metallicities (see also Fabbian et al. 2009a). After accounting for such corrections, Fabbian et al. (2009b) concluded that almost all diagnostics are now conceivably consistent, and that [O/Fe] exhibits a roughly flat plateau with values between +0.4 and +0.6 when [Fe/H] \( \lesssim -1\).

In contrast to the profusion of stellar studies of [O/Fe] in the VMP regime, this ratio has received relatively little attention in DLAs so far. The reason is that the most readily available O I absorption lines are almost always saturated, and the weakest lines are often blended with unrelated absorption in the Ly\( \alpha \) forest (Prochaska & Wolfe, 2002). For these reasons, several authors have investigated the [O/Fe] trend in sub-DLAs, where the O I absorption lines are weaker (Péroux et al., 2003b; O’Meara et al., 2005). However, uncertain negative ionisation corrections to the Fe II lines might become important for such systems, complicating the interpretation. Other authors have instead used [S/Zn] as a proxy for [O/Fe] (Nissen et al., 2007), but these lines become too weak in the VMP regime.

In fact, there have only been two studies in the literature that consider [O/Fe] in DLAs. The first was conducted by Petitjean, Ledoux, & Srianand (2008), who reported an [O/Fe] plateau of +0.32 \( \pm 0.10 \) from their sample of 13 DLAs with [Fe/H] < -1.0, (three of which have [Fe/H] < -2.0). The second, more recent, study was conducted by Penprase et al. (2010) whose sample includes five DLAs with [Fe/H] < -2.0. Their measurements, however, have large uncertainty (\( \sim \pm 0.4\) dex), so it is difficult to discern the underlying trend.

The survey presented here constitutes the largest sample of high resolution measures of O I and Fe II absorption in DLAs. The [O/Fe] values are plotted in Figure 4.11 where, for comparison, I also show a selection of [O/Fe] measurements in Galactic stars based on the forbidden [O I] \( \lambda 6300\) line (Nissen et al., 2002; Cayrel et al., 2004; García Pérez et al., 2006) with 3D corrections applied by Poul Nissen as I now describe.

The [O I] \( \lambda 6300\) line corresponds to a forbidden transition between two levels of the ground configuration of the O I atom, which are closely coupled via collisions. Because nearly all oxygen atoms are in the ground state in the atmospheres of late-type stars, one expects LTE to prevail; this is confirmed by detailed statistical equilibrium calculations (Kiselman, 1993). Non-LTE effects on the derived iron abundances are also negligible, when lines from the dominating ionisation stage (Fe II) are considered (Mashonkina et al., 2011). The 3D – 1D corrections are, however, significant for metal-poor stars; [O/H] derived from the [O I] \( \lambda 6300\) line decreases and [Fe/H] from Fe II lines increases slightly. As calculated by Nissen et al. (2002), the net effect on [O/Fe] for metal-poor main-sequence stars can be approximated by the expression
Figure 4.11: The [O/Fe] ratio in VMP DLAs (filled red triangles) where [O/H] has been measured from high resolution spectra. The green triangle refers to the carbon-enhanced DLA that will be considered separately in Chapter 5; the upper limit on [Fe/H] and corresponding lower limit on [O/Fe] (red triangle with arrows) is for the DLA towards J0311$-$1722 where Fe II absorption is not detected. The VMP DLA sample is compared with the stellar abundance measurements by Nissen et al. (2002) (circles), Cayrel et al. (2004) (squares) and García Pérez et al. (2006) (triangles), all based on the [O I] $\lambda$6300 line and corrected for 3D effects (see text for further details). The error bars at the top right corner of the plot indicate the typical errors in the stellar abundance measurements. All the measurements in this plot are tabulated, together with their individual errors, in Appendix C.

\[ [\text{O/Fe}]_{\text{3D}} - [\text{O/Fe}]_{\text{1D}} = 0.11 \text{[Fe/H]}, \quad \text{whereas} \quad [\text{Fe/H}]_{\text{3D}} - [\text{Fe/H}]_{\text{1D}} = -0.04 \text{[Fe/H]} \]. Cayrel et al. (2004) assumed that the same corrections are also valid for metal-poor red giants. Poul Nissen has recently verified that this is approximately correct by applying 3D corrections for giant stars as calculated by Collet, Asplund, & Trampedach (2007) for the [O I] $\lambda$6300 line and the Fe lines used by Cayrel et al. (2004). The same 3D corrections are then also expected for the cool subgiants studied by García Pérez et al. (2006), because they have atmospheric parameters intermediate between those of the main-sequence and red giant stars. For reference, I list in Appendix C values of [O/Fe] and [Fe/H] corrected for 3D effects (provided by Poul Nissen), as well as the [O/Fe] and [Fe/H] values measured in this sample of VMP DLAs.

Once the above-mentioned 3D corrections are applied, it can be seen that stars and DLAs share a similar trend of [O/Fe] with decreasing metallicity. To illustrate this point, I show the two samples (where [Fe/H] $\leq$ −2.0) as histograms in Figure 4.12. A K-S test\(^{10}\) reveals that the probability the two data sets are drawn from the same parent population is 71%.

The DLA values of [O/Fe] exhibit relatively little scatter given the errors. In the range −3 $\leq$ [Fe/H] $\leq$ −2, [O/Fe] is consistent with a constant value: \([\langle \text{O/Fe} \rangle] = +0.35 \pm 0.09\). This is

\(^{10}\)I have not included the star CS 22949$-$037 in this test (from Cayrel et al. 2004), since it exhibits a peculiar abundance pattern, with [O/Fe] = +1.54.
Figure 4.12: The distribution of $[O/\text{Fe}]$ values (where $[\text{Fe/H}] \leq -2.0$) in DLAs for this sample (red histogram) as compared to the compilation of $[O/\text{Fe}]$ in stars measured from the $[\text{O}I] \lambda 6300$ line (blue histogram). The dashed vertical red and blue lines indicate the median values for DLAs and stars respectively, corresponding to a difference of only $\sim 0.01$ dex.

In good agreement with the mean value reported by Petitjean, Ledoux, & Srianand (2008), \[ \langle [O/\text{Fe}] \rangle = +0.32 \pm 0.10, \text{ for } -2.0 \lesssim [\text{Fe/H}] \lesssim -1.0. \] Interestingly, these data suggest that there may be a hint that $[O/\text{Fe}]$ increases further when $[\text{Fe/H}] \lesssim -3$ (see Figure 4.11), but more data in this extremely metal-poor regime are required for firm conclusions. For the moment, I simply conclude that the ‘cosmic’ trend of $[O/\text{Fe}]$ in the VMP regime ($-3 \leq [\text{Fe/H}] \leq -2$) reaches a plateau of $\sim +0.35$, and is remarkably tight, especially given the observational errors.

In closing, the DLA measurements of $[O/\text{Fe}]$ help resolve the controversy regarding the relative abundances of O and Fe in metal-poor Galactic stars. It is plausible that VMP DLAs harbour the reservoir of neutral gas that will later condense to form a population of VMP stars. It is thus expected that towards the lowest metallicities both stars and DLAs should exhibit a similar trend. I conclude that, unless there are marked differences between the chemical evolution histories of DLAs and the early Galaxy, the results presented here and those of Petitjean, Ledoux, & Srianand (2008) favour an approximately constant plateau of stellar $[O/\text{Fe}]$ values when $[\text{Fe/H}] \lesssim -1$, with perhaps a mild increase with decreasing $[\text{Fe/H}]$. In any case, given the on-going improvement in the accuracy of the stellar models and the increasing samples of VMP DLAs, I anticipate that this issue may well be settled in the near future.

4.7 Discussion

4.7.1 The typical VMP DLA

With the large sample of measurements assembled here, one can now attempt to reconstruct the abundance pattern of a ‘typical’ VMP DLA, and to consider the clues it may provide on the nucleosynthesis by the earliest generation of stars. To achieve this goal, I am obviously required to select a reference element other than H, as there are no means to determine how much H was
The abundance pattern of a typical VMP DLA is illustrated by the black boxes, where the height in each box represents the dispersion in the population. The dashed line corresponds to the solar abundance ratios.

mixed with the nucleosynthetic products from the earliest generations of stars. Rather, the ratio of two metals provides the best handle for determining the properties of the generation of stars from which they were synthesised.

With the goal to probe early nucleosynthesis borne in mind, the most appropriate reference element is O, for the following reasons: (1) the dominant O yield comes from a single source – massive stars. Thus, the origin of O is well-understood; (2) O is the most abundant metal in the Universe; and (3) at the lowest metallicities, where one expects to uncover the signature from early nucleosynthesis, several O lines become unsaturated in DLAs. Thus, it is relatively straightforward to measure $[\text{O/H}]$. Taking oxygen then as the reference element, I have constructed the typical abundance pattern for a VMP DLA by determining the mean $\langle X/O \rangle$ ratio for each available element, $X$, and then referring this mean value to the adopted solar scale (Appendix A), i.e. I take the log of the mean, $\langle \log(X/O) \rangle$. These mean values are listed in Table 4.12 along with the dispersion in the available measurements ($\sigma_{\langle X/O \rangle}$). In the last column of Table 4.12, I indicate the total number of DLAs that were used in determining the mean $X/O$.

The corresponding abundance pattern is illustrated in Figure 4.13. In the following subsection, I investigate the most likely origin of the metals in VMP DLAs by directly comparing this typical abundance pattern to model yield calculations of both Population II and Population III stars. Before continuing, however, it is important to keep in mind that the ‘typical’ $\langle N/O \rangle$ in Table 4.12 may be biased high, because it does not include a number of upper limits, where the N I lines are too weak to be measured. As discussed above (Section 4.6.1), it is also possible that $\langle C/O \rangle$ may be biased towards lower values than the true mean.

<table>
<thead>
<tr>
<th>Element</th>
<th>$\langle X/O \rangle$</th>
<th>$\sigma_{\langle X/O \rangle}$</th>
<th>$n_{\langle X/O \rangle}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>C</td>
<td>-0.28</td>
<td>0.12</td>
<td>9</td>
</tr>
<tr>
<td>N</td>
<td>-1.04</td>
<td>0.23</td>
<td>13</td>
</tr>
<tr>
<td>Al</td>
<td>-0.43</td>
<td>0.14</td>
<td>11</td>
</tr>
<tr>
<td>Si</td>
<td>-0.08</td>
<td>0.10</td>
<td>21</td>
</tr>
<tr>
<td>Fe</td>
<td>-0.39</td>
<td>0.11</td>
<td>20</td>
</tr>
</tbody>
</table>
4.7 Discussion

4.7.2 Clues to early episodes of nucleosynthesis

There is increasing evidence to suggest that the most metal-poor DLAs may retain the signature from the earliest episodes of star formation (Erni et al., 2006; Pettini et al., 2008a). In this picture, the most metal-poor DLAs condensed directly out of material that was enriched by either: (1) an external halo that distributed its products over large cosmological volumes via multi-SN events (Madau, Ferrara, & Rees, 2001), or perhaps (2) just a few SNe from the halo in which the DLA now resides (Bland-Hawthorn et al., 2011).

Cosmological simulations of galaxy formation support the possibility that such DLAs still retain the chemical signature of early enrichment (Pontzen et al., 2008; Tescari et al., 2009); the most metal-poor DLAs arise in low mass halos that have undergone little to no in situ star formation. It is possible, however, that VMP DLAs acquired some of their metals at later times from nearby sources that delivered metals into the IGM via galactic superwinds (see e.g. Oppenheimer & Davé 2008), which may complicate the interpretation. Perhaps the most straightforward way to discriminate between these enrichment scenarios is to compare the model yields of both Population III and Population II stars with that of the typical VMP DLA described in Section 4.7.1.

I consider three sources that could be responsible for the metals in VMP DLAs: (1) massive metal-free stars, with main sequence masses in the range 140 – 260 M⊙ that explode as pair-instability SNe (PISN; Heger & Woosley 2002); (2) massive metal-free stars, with progenitor masses in the range 10 – 100 M⊙ that explode as core-collapse SNe (CCSN; Heger & Woosley 2010); and (3) massive Population II (and I) stars, with progenitor masses in the range 13 – 35 M⊙, covering a range in metallicity (Chieffi & Limongi, 2004), also ending their lives as CCSN.

To determine the dominant source of the metals in VMP DLAs, I have integrated these model yields over a Salpeter-like power law IMF, \( dN/dM \propto M^{-(1+\gamma)} \) (where \( \gamma = 1.35 \) for a Salpeter IMF), and consider three values for the power law index in the case of zero metallicity (\( \gamma = 1.35, 2.35, 3.35 \)). For massive Population II stars, I consider only a Salpeter IMF, with \( \gamma = 1.35 \). The results of these calculations are shown in Figure 4.14.

Consider first the PISNe, the yields of which are the least model dependent of the three models considered here. Qualitatively, these models could have been ruled out on the basis of the near-solar [Si/O] that is typical of the VMP DLA population; PISN are expected to yield supersolar [Si, S, Ar, Ca/O]. Indeed, for the range of \( \gamma \) considered here, such SNe provide a poor fit to the typical VMP DLA population (top panel of Figure 4.14).

I now turn to models of massive stars (\( \gtrsim 10 \text{ M}_\odot \)) that end their lives as CCSNe. Unfortunately, the explosion mechanism of CCSN is poorly understood, and several unknown physical effects need to be parameterised and suitably adjusted to find the best solution for a given set of data. In particular, one usually parameterises the explosion energy, the degree of mixing between the stellar layers during the explosion, and the amount of material that falls back onto the central remnant.
Figure 4.14: The abundance pattern of the ‘typical’ VMP DLA (black boxes; cf. Figure 4.13) is compared to nucleosynthesis models of massive stars (symbols connected by lines). Top panel: The symbols represent the yields from pair-instability supernovae (PISNe) of zero metallicity stars (Heger & Woosley, 2002) for three indices of a power law IMF (red, green, blue corresponds to $\gamma = 1.35, 2.35, 3.35$ respectively). Middle panel: Same as above, except for core-collapse supernovae (CCSNe) models of zero metallicity stars (Heger & Woosley, 2010). Bottom panel: Comparing the typical DLA abundance pattern with the model explosive yields of massive Population II stars (Chieffi & Limongi, 2004). A range of metallicities is considered, as indicated by the accompanying legend. In all cases for this bottom panel, I have plotted the integrated yields from the Salpeter IMF with a power law index of $\gamma = 1.35$. 

![Graph showing abundance pattern comparison](image-url)
The most recent suite of published CCSN yields for metal-free stars are those by Heger & Woosley (2010). These computations provide a detailed account of the nucleosynthetic products over a large range of progenitor masses (10 – 100 M⊙) with a typical mass resolution of 0.1 M⊙. As part of their study, Heger & Woosley (2010) compiled a database of these model yields that are imported into their STARFIT software. This software is designed to objectively sieve through the vast parameter space and select the explosion parameters that best fit the data.

To maintain consistency with the other yield models that are considered here, I have compared the abundance pattern of the typical VMP DLA to the expected yields for the three power law indices of a Salpeter-like IMF (γ = 1.35, 2.35, 3.35). I therefore froze the remaining parameters to the ‘standard’ case, which corresponds to a constant explosion energy (1.2 × 10^51 erg) for all masses in the range 10 – 100 M⊙ (see Heger & Woosley 2010 for further details). This standard case sets the piston location of the explosion to be at the base of the oxygen burning shell (where the entropy per baryon ≃ 4), and applies a mixing boxcar filter with a width which is 10% the He core size. The material that falls back onto the central remnant is not parameterised in this code, but is instead calculated by switching the piston off 100 s after the explosion and defining an inner boundary condition where material is accreted.

All three fits are reproduced in the middle panel of Figure 4.14. The ‘standard’ case with the yields from metal-free stars seems to produce a reasonable agreement with the observed metal ratios in the ‘typical’ VMP DLA, although Al is discrepant by ~ 0.5 dex and, interestingly, IMF slopes steeper than Salpeter seem to fit the data best. Of course, by relaxing some of the default constraints it may be possible to improve the fit further, but I have refrained from doing so, given that there are still many uncertainties in accurately modelling the physics behind the explosion. Some of these uncertainties are only now beginning to be addressed in some detail (see Joggerst, Almgren, & Woosley 2010b and references therein).

Finally, I consider the set of model CCSN yields published by Chieffi & Limongi (2004), which allow me to test whether or not Population II stars can also account for the origin of the metals in VMP DLAs. Before comparing the models by Chieffi & Limongi to the typical VMP DLA, it is worth noting the important differences between this code and the one described above by Heger & Woosley (2010). Aside from the obvious difference in metallicity, the models by Chieffi & Limongi (2004) target the mass range 13 – 35 M⊙ with a relatively coarser mass resolution of 5 M⊙. In addition, this code parameterises the amount of material that falls back onto the central remnant. In their standard case, this prescription requires 0.1 M⊙ of 56Ni to be ejected from the star, and thus all material interior to this mass coordinate is ‘accreted’ by the remnant. Finally, this code is yet to implement a ‘mixing parameter’ to model the mixing that takes place between the stellar layers during the SN explosion. Adopting their standard case, which has an explosion energy of 1.2 × 10^51 erg, I have integrated the model yields over a Salpeter-like IMF with γ = 1.35. The results are shown in the bottom panel of Figure 4.14.

11STARFIT is written with the INTERACTIVE DATA LANGUAGE software and is available from: http://homepages.spa.umn.edu/~alex/znuc/
These calculations adopt an initial chemical composition that is simply scaled from the solar abundance pattern. According to Chieffi & Limongi (2004), the model yields for stars with an initial metallicity of $Z \leq 0.005 Z_{\odot}$ are not strongly dependent on the initial composition of the star. It is important to note, however, that by introducing an $\alpha$-enhancement to the initial metallicity (which is perhaps more realistic than simply scaling the solar abundance pattern), the yields for the odd atomic number elements are increased (see their Figure 1). Thus, these calculations may underestimate the N/O and Al/O ratios. Furthermore, at higher metallicities ($Z > 0.005 Z_{\odot}$), the yields do depend on the initial composition of the stars.

Inspecting the bottom panel of Figure 4.14, it can be seen that, at face value, the $Z/Z_{\odot} = 0.0, 0.05$ and 0.3 models provide reasonable fits to the abundance pattern of a typical VMP DLA. I also note the broad agreement between the metal-free models by Heger & Woosley (2010) and Chieffi & Limongi (2004). On the other hand, the $Z/Z_{\odot} = 0.05$ and 0.3 models also seem to provide reasonable fits to the abundance pattern. However, once the stars have reached metallicities greater than 1/20 of solar, the metal ratios in the gas should be compared with the predictions of full chemical evolution models which are beyond the scope of this thesis.

Given the current (largely) model-dependent nature of these calculations, I am unable to draw firm conclusions at this stage. Whilst the above models suggest that metal-free stars could have synthesised the metals that now reside in VMP DLAs, I cannot rule out the possibility that Population II stars are also responsible. I suspect that it will be necessary to measure the abundances of additional metals in order to better distinguish between Population II and Population III models, because the ratios of the most abundant elements ([C/O], [Si/O] and [Fe/O]) provide the weakest constraints on the nature of the objects that synthesised them.

Finally, I note that the models used here to compare with a typical VMP DLA are still quite dependent on unknown physics; the single largest uncertainty in these models is the explosion mechanism. Additional physics also needs to be included, such as the mixing induced by stellar rotation (Meynet, Ekström, & Maeder, 2006; Hirschi, 2007; Meynet et al., 2010; Joggerst et al., 2010a) and the Rayleigh-Taylor instability (Joggerst, Woosley, & Heger, 2009). These effects will presumably be considered in the next generation of fine-grid nucleosynthesis models, when the limitations of computing power will hopefully be less of a concern.

### 4.7.3 Comparison with data of medium spectral resolution

Finally, I compare DLA abundance determinations obtained from high ($R \sim 40000$) and medium ($R \sim 5000$) spectral resolution data. Such a comparison is motivated by the realisation that, even with efficient echelle spectrographs on 8–10 m telescopes, it is typically necessary to integrate on a single QSO for the equivalent of one night in order to obtain the S/N ratio required to measure elemental abundances from high resolution spectra. By settling for lower resolutions, the exposure times are greatly reduced; for example, most of the QSOs in the survey by Penprase et al. (2010) were observed for about one hour with the Echellette Spectrograph and Imager.
Table 4.13: Column densities estimated from high and medium spectral resolution spectra

<table>
<thead>
<tr>
<th>Ion</th>
<th>( \log N_{\text{c,oo}}^a )</th>
<th>( \log N_{\text{pen,m}}^b )</th>
<th>( \log N_{\text{pen,c}}^c )</th>
</tr>
</thead>
<tbody>
<tr>
<td>J0831+3358</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>Al II</td>
<td>12.19 ± 0.06</td>
<td>...</td>
<td>&gt; 11.56</td>
</tr>
<tr>
<td>Si II</td>
<td>13.75 ± 0.04</td>
<td>...</td>
<td>&gt; 13.03</td>
</tr>
<tr>
<td>Fe II</td>
<td>13.33 ± 0.06</td>
<td>...</td>
<td>&gt; 13.13</td>
</tr>
<tr>
<td>J1001+0343</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>C II</td>
<td>13.58 ± 0.02</td>
<td>13.63 ± 0.06</td>
<td>13.76 ± 0.13</td>
</tr>
<tr>
<td>O I</td>
<td>14.25 ± 0.02</td>
<td>13.90 ± 0.07</td>
<td>13.98 ± 0.08</td>
</tr>
<tr>
<td>Si II</td>
<td>12.86 ± 0.01</td>
<td>12.70 ± 0.05</td>
<td>12.82 ± 0.12</td>
</tr>
<tr>
<td>J1037+0139</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>O I</td>
<td>15.06 ± 0.04</td>
<td>...</td>
<td>&gt; 14.48</td>
</tr>
<tr>
<td>Al II</td>
<td>12.32 ± 0.03</td>
<td>...</td>
<td>&gt; 12.09</td>
</tr>
<tr>
<td>Si II</td>
<td>13.97 ± 0.03</td>
<td>...</td>
<td>&gt; 13.69</td>
</tr>
<tr>
<td>Fe II</td>
<td>13.53 ± 0.02</td>
<td>...</td>
<td>&gt; 13.33</td>
</tr>
</tbody>
</table>

\(^a\)My results.
\(^b\)Column densities measured by Penprase et al. (2010).
\(^c\)Column density estimates by Penprase et al. (2010) after applying a saturation correction.

(mounted on the Keck II telescope). Clearly, it is of interest to test how similar the abundance measurements are for the most metal-poor DLAs, if one were to forgo the accuracy of high spectral resolution in order to secure a larger sample.

In this context, there are two main concerns that potentially limit the accuracy of abundance measurements from medium (as opposed to high) spectral resolution data. First, VMP DLAs typically have line widths \(< 10 \text{ km s}^{-1}\) (Ledoux et al., 2006; Murphy et al., 2007; Prochaska et al., 2008), which are unresolved at \( R = 5000 \). One must therefore appeal to a curve-of-growth analysis appropriate to a single absorbing cloud, which is often an oversimplification (see Table 4.2, and also the discussion by Prochaska 2006). Second, the relevant absorption lines may be saturated, and the degree of saturation may be difficult to estimate correctly, even with a well-defined curve-of-growth.

As it happens, three VMP DLAs from the present work are in common with the lower resolution survey of Penprase et al. (2010), providing me with the means to compare column densities from the two sets of spectra, as in Table 4.13. For two of the DLAs, J0831+3358 and J1037+0139, Penprase et al. (2010) reported lower limits on the column densities of the available metal ions. As can be seen from Table 4.13, while these lower limits are always consistent with the values measured from the high resolution echelle spectra, they fall short of the true column density by widely differing amounts, from as little as \(-0.13\) dex to as much as \(-0.72\) dex. This wide range significantly reduces the usefulness of the lower limits.

Turning now to the DLA at \( z_{\text{abs}} = 3.07841 \) towards J1001+0343, I recall that the UVES
spectra show that the absorption arises in a single component with Doppler parameter $b = 7.0 \pm 0.1 \text{ km s}^{-1}$ (see Figure 4.4). This value is not too dissimilar from $b = 7.5 \text{ km s}^{-1}$ estimated by Penprase et al. (2010) from a curve-of-growth analysis. Indeed, the two analyses give consistent estimates of the Si \text{II} column density (after Penprase et al. (2010) apply a saturation correction). The C \text{II} and O \text{I} lines, however, tell a different story. For this DLA, Penprase et al. (2010) need not have applied a saturation correction to the C \text{II} $\lambda 1334$ line, since it is not strongly saturated (see Figure 4.4). Indeed, prior to applying such a correction, their column density estimate was in broad agreement with that derived here. Conversely, O \text{I} $\lambda 1302$ is closer to saturation and, even with the correction applied by Penprase et al. (2010), these authors’ estimate falls short of the value deduced here by nearly a factor of two. The combined effect is an overestimate of [C/O] by 0.45 dex.

Based on this example, there appear to be non-negligible uncertainties in the derivation of element ratios from spectra at $R \sim 5000$. While these uncertainties did not affect the principle goal of the study by Penprase et al. (2010) – to uncover the most metal-poor DLAs – it would appear that high resolution observations are indeed necessary to measure element abundances in VMP DLAs with an accuracy better than a factor of $\sim 2$.

### 4.8 Summary and Conclusions

I have conducted a survey for very metal-poor DLAs to shed light on the earliest episodes of nucleosynthesis in our Universe. This sample includes seven new DLAs observed with high resolution spectrographs ($R \gtrsim 30000$); when combined with the other metal-poor DLAs previously reported from this programme (Pettini et al., 2008a) and an additional ten DLAs from the literature, it constitutes the largest survey to date for DLAs with a metallicity $[\text{Fe/H}] < -2.0$.

From the analysis of these data, I draw the following conclusions.

(i) Having now doubled the sample of DLAs where the C/O ratio is measured from unsaturated absorption lines, I confirm that DLAs exhibit near-solar values of C/O at the lowest metallicities probed. Furthermore, I find good agreement in the C/O ratio observed in this sample of DLAs and in recent compilations of the most metal-poor Galactic halo stars. I argue that such good agreement points to a universal origin for the C/O ‘excess’ in this regime.

(ii) For the first time, I have investigated the [O/Fe] ratio in very metal-poor DLAs. For 20 DLAs with $[\text{Fe/H}] < -2.0$, I find a small dispersion around a mean value $\langle [\text{O/Fe}] \rangle = +0.39 \pm 0.11$. I have also presented tentative evidence for a rise in the [O/Fe] ratio when $[\text{Fe/H}] \lesssim -3.0$.

(iii) In view of the long-standing debate as to the behaviour of the [O/Fe] ratio in metal-poor Galactic halo stars, I have compared the stellar trend to that observed in this sample of DLAs. I find good agreement between stars and DLAs when the stellar oxygen abundance is measured from the [O \text{I}] $\lambda 6300$ line (after correcting for 3D effects). Based on the available DLA samples,
I conclude that \( [\text{O}/\text{Fe}] \) is essentially flat in the metallicity interval \(-3.0 \lesssim [\text{Fe/H}] \lesssim -1.0\), with the possibility of an increase at yet lower metallicities.

(iv) I have constructed the abundance pattern of a typical very metal-poor DLA for the five most commonly observed metals, using O as a reference. I find that Si/O is just below solar \( \langle [\text{Si}/\text{O}] \rangle = -0.08 \), whilst \( \langle [\text{C}/\text{O}] \rangle = -0.28 \) and \( \langle [\text{Fe}/\text{O}] \rangle = -0.39 \). The largest deviations from a solar scaled abundance pattern are exhibited by N and Al, with \( \langle [\text{N,Al}/\text{O}] \rangle = -1.04, -0.43 \).

(v) One of the main aims of this work was to investigate the origin of the metals in the most metal-poor DLAs. To achieve this goal, I compared the abundance pattern of a ‘typical’ VMP DLA with those expected from model calculations using the yields of Population II and Population III stars. For the few elements considered here, I find a reasonable agreement between the abundance pattern of the typical VMP DLA and the ‘standard model’ of a population of metal-free stars (i.e. a top-heavy initial mass function where all stars explode as core-collapse supernovae with an energy of \( 1.2 \times 10^{51} \) erg). However, given that I only have access to a handful of metals, I cannot unambiguously rule out (an additional contribution from) more metal-rich Population II stars. On the other hand, I am able to firmly conclude that the typical very metal-poor DLA was not solely enriched by pair-instability supernovae from very massive metal-free stars.

This ongoing programme to measure the abundances in the most metal-poor DLAs complements local studies of Galactic metal-poor halo stars. The good agreement I have found between these two populations suggests a universal origin for their metals. The results presented here emphasise the importance of measuring elemental abundances in the most metal-poor DLAs; these systems present us with a unique window of opportunity to probe the nucleosynthesis by some of the earliest structures in the Universe.
The nucleosynthesis from the first stars

Chapter Summary

I present high resolution observations of an extremely metal-poor damped Ly$\alpha$ system, at $z_{\text{abs}} = 2.3400972$ in the spectrum of the QSO J0035−0918, exhibiting an abundance pattern consistent with model predictions for the supernova yields of Population III stars. Specifically, this DLA has $[\text{Fe/H}] \simeq -3$, shows a clear ‘odd-even’ effect, and is C-rich with $[\text{C/Fe}] = +1.53$ (with a strict lower limit of $[\text{C/Fe}] \geq +0.46$), a factor of $\sim 20$ greater than reported in any other damped Ly$\alpha$ system. In analogy to the carbon-enhanced metal-poor stars in the Galactic halo (with $[\text{C/Fe}] > +1.0$), this is the first reported case of a carbon-enhanced damped Ly$\alpha$ system. I determine an upper limit to the mass of $^{12}\text{C}$, $M(^{12}\text{C}) \leq 200\, M_\odot$, which depends on the unknown gas density $n(\text{H})$; if $n(\text{H}) > 1\, \text{cm}^{-3}$ (which is quite likely for this DLA given its low velocity dispersion), then $M(^{12}\text{C}) \leq 2\, M_\odot$, consistent with pollution by only a few prior supernovae. I speculate that DLAs such as the one discussed in this Chapter may represent the ‘missing link’ between the yields of Pop III stars and their later incorporation in the class of carbon-enhanced metal-poor stars which show no enhancement of neutron-capture elements (CEMP-no stars).
Chapter 5. The nucleosynthesis from the first stars

5.1 Introduction

DLAs are the neutral gas reservoirs at high redshift that have, by definition, neutral hydrogen column densities in excess of $10^{20.3}$ atoms cm$^{-2}$ (see Wolfe et al. 2005 for a review). At these high column densities the gas is self-shielded (e.g. Vladilo et al., 2001), resulting in a simple ionisation structure which facilitates the derivation of element abundances. Moreover, the abundances thus derived are independent of the geometrical configuration and thermodynamical state of the gas, and of most other factors which complicate the analysis of stellar spectra (e.g. Asplund, 2005). The largest uncertainties in DLA abundance studies are due to the effects of line saturation and dust depletion (for some elements), although the latter of these concerns is found to be minimal when the metallicity of the DLA is below $1/100Z_\odot$ (Pettini et al., 1997a; Prochaska & Wolfe, 2002; Akerman et al., 2005).

In recent years, these very metal-poor (VMP) DLAs, with metallicities [Fe/H] $<-2$, have attracted an increasing amount of interest because of their potential for probing gas which may still bear the chemical imprint of the first few generations of stars to have formed in the Universe (e.g. Erni et al., 2006). They are thus an extremely valuable complement, at high redshifts, to local studies of the oldest and most metal-poor stars in the Galactic halo.

The VMP DLA regime was largely unexplored until very recently, when it became possible to identify candidate metal-poor DLAs in Sloan Digital Sky Survey (SDSS) quasars, and then measure their chemical composition with high resolution follow-up spectroscopy. The first high spectral resolution (R $\gg 30000$, FWHM $\lesssim 10$ km s$^{-1}$) sample of VMP DLAs was compiled by Pettini et al. (2008a), whose main goal was to compare the relative abundances of C, N, and O with the trends observed in VMP halo stars in our Galaxy. More recently, Penprase et al. (2010) have presented medium resolution (FWHM $\sim 60$ km s$^{-1}$) spectroscopy of a sample of 27 VMP DLAs to explore the general properties of DLAs in the VMP regime. However, as Penprase et al. acknowledge themselves, there are difficulties with accurately measuring column densities from medium (as opposed to high) resolution data, given that most DLAs with metallicities less than $1/100$ solar exhibit very low velocity dispersions, with metal line widths less than 10 km s$^{-1}$ (Ledoux et al., 2006; Murphy et al., 2007; Prochaska et al., 2008). Under these circumstances, line saturation can easily be overlooked.

Perhaps the most startling result from abundance studies of VMP DLAs are the near-solar values of [C/O] at low metallicity (Pettini et al., 2008a), in line with the peculiar upturn in the [C/O] abundance below [O/H] $\sim -1.0$ in Galactic halo stars discovered by Akerman et al. (2004) and later confirmed by Fabbian et al. (2009a). Akerman et al. (2004) attributed this behaviour to an increased C yield from earlier generations of massive stars. Penprase et al. (2010) extended this work and reported a number of DLAs with supersolar [C/O] and suggested that this ratio continues to increase with decreasing [O/H]. The more recent compilation of high spectral resolution observations of VMP DLAs discussed in Chapter 4, however, suggests that the ratio in fact plateaus at [C/O] $\sim -0.2$. 
In contrast to the relatively new interest in VMP DLAs, metal-poor studies of Galactic halo stars have received ongoing attention for more than two decades, most recently from the dedicated HK (Beers, Preston, & Shectman, 1992) and HES (Christlieb et al., 2001) surveys. A relevant result emerging from this work is that nearly one-quarter of all metal-poor stars with \[\text{[Fe/H]} < -2.0\] exhibit a marked carbon enhancement, with \[\text{[C/Fe]} > +1.0\] (Beers & Christlieb, 2005; Lucatello et al., 2006). These are collectively known as carbon-enhanced metal-poor stars (CEMP stars), and have been subdivided into four classes based on the abundances of their neutron-capture elements: (i & ii) The CEMP-s and CEMP-r classes, with enhancements of elements produced predominantly by the \(s\)-process and \(r\)-process respectively; (iii) the CEMP-rs class, with enhancements in both the \(s\)- and \(r\)-process elements; and (iv) the CEMP-no class, which exhibits no such enhancements.

For further details of these classes, and the likely origins of their carbon enhancements, I direct the reader to Masseron et al. (2010). In short, there is reasonable evidence to suggest that CEMP-s and CEMP-rs stars are extrinsically polluted by a now extinct asymptotic giant branch (AGB) companion. The origin of the CEMP-no class, however, is not yet firmly established. Whilst several models invoking mass transfer from an AGB companion could explain the lack of neutron-capture elements (e.g. Fujimoto, Ikeda, & Iben 2000; Siess, Goriely, & Langer 2004), radial-velocity measurements have not yet confirmed whether CEMP-no stars are host to binary companions. Indeed, the apparent difference in the metallicity distributions between CEMP-no stars and the other CEMP classes, in the sense that CEMP-no stars are more numerous at lower metallicity, might suggest that a mechanism other than the AGB binary transfer scenario produces the C enhancement in CEMP-no stars (Aoki et al., 2007). In all likelihood, as pointed out by Masseron et al. (2010), there is a continuous link that connects some CEMP-no stars with CEMP-s stars, while the carbon enhancement of other CEMP-no stars may have a different origin.

Models of core-collapse supernova yields from Population III (or near metal-free) stars do entertain high C yields relative to Fe (e.g. Woosley & Weaver 1995). Population III enrichment is a particularly intriguing explanation for the origin of the carbon enhancement in some CEMP stars, since the three most Fe-poor halo stars all exhibit carbon enhancements relative to iron (Christlieb et al., 2002; Frebel et al., 2005; Norris et al., 2007, see Figure 1.9). Moreover, the fraction of metal-poor stars that exhibit a carbon-enhancement increases with decreasing metallicity (Beers & Christlieb, 2005). In this Population III enrichment scenario, the carbon enhancement in the extremely metal-poor regime reflects the initial composition of the gas from which these stars formed, rather than resulting from mass transfer from an evolved companion.

Whilst the physics behind core-collapse supernovae is poorly constrained, several parameterised models have been developed to calculate the expected yields from zero-metallicity progenitors to compare with the observations of the most Fe-poor CEMP stars, as well as the more ‘normal’ non-CEMP stars. The most important (and largely unknown) quantities are the degree of mixing during the supernova explosion and the amount of fallback onto the rem-
nant black hole thereafter (Umeda & Nomoto, 2003). Umeda & Nomoto (2003) parameterised both quantities, then suitably selected the appropriate values that reproduce the observed stellar abundance patterns. Heger & Woosley (2010), on the other hand, parameterise only the mixing parameter by applying a running boxcar filter over the star following the explosion. This prescription successfully reproduces the chemical composition of the extremely metal-poor (non-CEMP) halo stars from the study by Cayrel et al. (2004), as well as that of the most Fe-poor CEMP stars HE0107−5240 (Christlieb et al., 2002) and HE1327−2326 (Frebel et al., 2005). Joggerst, Woosley, & Heger (2009) extended this work by mapping the one-dimensional models by Heger & Woosley (2010) onto a two-dimensional grid to follow Rayleigh-Taylor induced mixing after explosive nuclear burning. Whilst the Joggerst, Woosley, & Heger (2009) models are physically motivated, they are unable to reproduce the relatively high levels of nitrogen enrichment that are observed in the most Fe-poor stars, nor are they able to produce sufficient Fe-peak elements.

These concerns are alleviated with models that include rotation (e.g. Meynet, Ekström, & Maeder 2006; Hirschi 2007), since rotation induces additional mixing and mass loss in the presupernova phase of low-metallicity stars. The only simulation available to date that investigates the effects of both rotational and Rayleigh-Taylor mixing, as well as incorporating a realistic prescription of fallback, are those presented recently by Joggerst et al. (2010a) (see also Joggerst, Almgren, & Woosley 2010b). By introducing rotation in their zero-metallicity models, Joggerst et al. (2010a) report an increased nitrogen yield, as well as an increased Fe-peak element yield, however, these models are not able to reproduce the high CNO to Fe ratios observed in the most metal-poor stars. In summary, it is still a matter of some debate whether these extremely metal-poor early Population II stars: (i) were born out of gas which had previously been carbon-enriched by the first stars; or (ii) received a CNO top-up from a companion star; and/or (iii) exhibit a degree of self-pollution from their own nucleosynthesis.

The first possibility, that at least some CEMP stars were born out of carbon-enhanced gas, is given additional support from the discovery reported here of an extremely metal-poor DLA at $z_{\text{abs}} = 2.3400972$ exhibiting a carbon enhancement akin to that measured in Galactic CEMP stars. No other case was known until now, given (i) the rarity of DLAs with $[\text{Fe/H}] < -2$, which lie in the tail of the metallicity distribution of DLAs (e.g. Pontzen et al., 2008), and (ii) the difficulty in measuring $[\text{C/H}]$ in DLAs, where the relevant absorption lines are often saturated even in the VMP regime. I speculate that this DLA may be the ‘missing link’ between the first few generations of stars and some of the CEMP stars in the Galactic halo.

This Chapter is organised as follows. Section 5.2 summarises the observations and data reduction. I analyse the absorption lines in the damped Ly$\alpha$ system and deduce corresponding element abundances in Section 5.3. In Section 5.4 I scrutinise potential issues that could masquerade as a carbon enhancement. After ruling out these possibilities, I discuss in Section 5.5 possible origins of this enhancement, and consider the implications of this finding for models of CEMP stars. Finally, I summarise the results and draw my conclusions in Section 5.6.
5.2 Observations and Data Reduction

The $m_r = 18.89$, $z_{\text{em}} = 2.42$ QSO J0035−0918 was identified as a candidate VMP DLA for the survey presented in Chapter 4 on the basis of its SDSS spectrum which shows a DLA at $z_{\text{abs}} \simeq 2.340$ with no apparent associated metal lines. Such cases are highly suggestive of narrow and weak metal absorption lines which are undetectable at the coarse resolution and signal-to-noise ratio (S/N) of the discovery SDSS spectra.

Follow-up observations of J0035−0918 were made with the Magellan Echellette (MagE) spectrograph (Marshall et al., 2008) on the Magellan II Clay telescope on the nights of 2008 December 30 and 31 in good conditions with 1 arcsec seeing. Two 2400 s exposures were taken using the standard setup with $1 \times 1$ binning and a 1.0 arcsec slit giving a resolving power $R \equiv \lambda/\Delta \lambda \simeq 4100$. Standard calibrations were taken following the recommendations by Simcoe, Hennawi & Williams\(^1\). The data were reduced by Regina Jorgenson using a custom set of IDL routines written by G. D. Becker as described in Becker et al. (2006). The MagE spectrum confirmed the VMP DLA nature of the $z_{\text{abs}} = 2.340972$ absorption system, by showing clear damping wings to the Ly$\alpha$ absorption line (see top panel of Figure 5.1) and unusually weak associated metal lines.

Encouraged by these initial indications, J0035−0918 was subsequently observed on the night of 2009 December 9 with the High Resolution Echelle Spectrograph (Vogt et al., 1994) on the Keck I telescope under good conditions with sub-arcsecond seeing, for a total integration time of 16200 s, divided into 6 exposures of 2700 s, resulting in a signal-to-noise ratio near 4500 at $S/N \simeq 18$. A 1.148 arcsec wide slit was used which, when uniformly illuminated, provides a nominal spectral resolution $R \equiv \lambda/\Delta \lambda = 37000$. From these spectra, I measure $R \simeq 41000$ which corresponds to a velocity full-width at half maximum FWHM = 7.3 km s$^{-1}$, sampled by $\sim 3$ pixels. I employed the UV cross-disperser which covers the wavelength range 3100−6000 Å with $\sim 70$ Å-wide gaps near 4000 Å and 5000 Å.

The HIRES spectra were reduced with the MAKEE data reduction pipeline developed by Tom Barlow, which performs the usual steps of flat-fielding, order tracing, background subtraction, and extraction of the final 1-D spectrum. The data were wavelength calibrated by reference to the spectrum of a ThAr lamp, and mapped onto a vacuum heliocentric wavelength scale. The extracted spectra were merged and then normalised by dividing out the QSO continuum and emission lines, using the software package UVES POPLER, maintained by Michael Murphy\(^2\). Following this step, all available metal absorption lines associated with the DLA were extracted in $\pm 150$ km s$^{-1}$ windows around the pixel with highest optical depth. A further fine correction to the continuum was then applied if necessary.

\(^1\)see http://web.mit.edu/rsimcoe/www/MagE/mage.html
\(^2\)UVES POPLER is available from http://astronomy.swin.edu.au/~mmurphy/UVES_pople
Table 5.1: Metal lines in the \( z_{\text{abs}} = 2.3400972 \) DLA towards J0035−0918

<table>
<thead>
<tr>
<th>Ion</th>
<th>Wavelength(^a) (Å)</th>
<th>( f^a )</th>
<th>( W_0^b ) (mÅ)</th>
<th>( \delta W_0^c ) (mÅ)</th>
<th>( \delta W_{0,\text{cont}}^d ) (mÅ)</th>
</tr>
</thead>
<tbody>
<tr>
<td>C II</td>
<td>1036.3367</td>
<td>0.118</td>
<td>39</td>
<td>2</td>
<td>1</td>
</tr>
<tr>
<td>C II</td>
<td>1334.5323</td>
<td>0.127</td>
<td>54</td>
<td>2</td>
<td>1</td>
</tr>
<tr>
<td>N I</td>
<td>1134.1653</td>
<td>0.0146</td>
<td>&lt; 5.3(^e)</td>
<td>...</td>
<td>...</td>
</tr>
<tr>
<td>N I</td>
<td>1134.4149</td>
<td>0.0278</td>
<td>6.2</td>
<td>1.7</td>
<td>0.5</td>
</tr>
<tr>
<td>N I</td>
<td>1134.9803</td>
<td>0.0416</td>
<td>12.5</td>
<td>1.6</td>
<td>0.5</td>
</tr>
<tr>
<td>N II</td>
<td>1083.9937</td>
<td>0.111</td>
<td>&lt; 4.6(^e)</td>
<td>...</td>
<td>...</td>
</tr>
<tr>
<td>O I</td>
<td>971.7382</td>
<td>0.0116</td>
<td>24</td>
<td>3</td>
<td>1</td>
</tr>
<tr>
<td>O I</td>
<td>988.5778</td>
<td>0.000553</td>
<td>&lt; 6.0(^e)</td>
<td>...</td>
<td>...</td>
</tr>
<tr>
<td>O I</td>
<td>988.6549</td>
<td>0.0083</td>
<td>23</td>
<td>2</td>
<td>1</td>
</tr>
<tr>
<td>O I</td>
<td>988.7734</td>
<td>0.0465</td>
<td>38</td>
<td>3</td>
<td>1</td>
</tr>
<tr>
<td>O I</td>
<td>1039.2304</td>
<td>0.00907</td>
<td>23</td>
<td>2</td>
<td>1</td>
</tr>
<tr>
<td>O I</td>
<td>1302.1685</td>
<td>0.048</td>
<td>42</td>
<td>2</td>
<td>1</td>
</tr>
<tr>
<td>Al II</td>
<td>1670.7886</td>
<td>1.740</td>
<td>15</td>
<td>2</td>
<td>0.5</td>
</tr>
<tr>
<td>Si II</td>
<td>989.8731</td>
<td>0.171</td>
<td>15</td>
<td>2</td>
<td>1</td>
</tr>
<tr>
<td>Si II</td>
<td>1193.2897</td>
<td>0.582</td>
<td>28</td>
<td>3</td>
<td>2</td>
</tr>
<tr>
<td>Si II</td>
<td>1260.4221</td>
<td>1.18</td>
<td>39</td>
<td>2</td>
<td>0.5</td>
</tr>
<tr>
<td>Si II</td>
<td>1304.3702</td>
<td>0.0863</td>
<td>21</td>
<td>2</td>
<td>1</td>
</tr>
<tr>
<td>Si II</td>
<td>1526.7070</td>
<td>0.133</td>
<td>34</td>
<td>2</td>
<td>1</td>
</tr>
<tr>
<td>S II</td>
<td>1253.805</td>
<td>0.0109</td>
<td>&lt; 2.0(^e)</td>
<td>...</td>
<td>...</td>
</tr>
<tr>
<td>S II</td>
<td>1259.5180</td>
<td>0.0166</td>
<td>3.7</td>
<td>0.8</td>
<td>0.3</td>
</tr>
<tr>
<td>Fe II</td>
<td>1260.533</td>
<td>0.0240</td>
<td>&lt; 2.5(^e)</td>
<td>...</td>
<td>...</td>
</tr>
<tr>
<td>Fe II</td>
<td>1608.4509</td>
<td>0.0577</td>
<td>10</td>
<td>2</td>
<td>0.5</td>
</tr>
</tbody>
</table>

\(^a\) Laboratory wavelengths and \( f \)-values from Morton (2003).
\(^b\) Rest-frame equivalent width.
\(^c\) Random error on the equivalent width \( W_0 \).
\(^d\) Systematic error on \( W_0 \) from 1\( \sigma \) uncertainty in the continuum placement.
\(^e\) 3\( \sigma \) upper limit on the rest-frame equivalent width.

5.3 Profile Fitting and Abundance Analysis

Table 5.1 lists all the metal absorption lines in the \( z_{\text{abs}} = 2.3400972 \) DLA detected in the HIRES spectrum of J0035−0918, together with a few interesting transitions which are below the detection limit of the data. For each line, I give the rest-frame wavelength and oscillator strength that I adopted for this work, as well as the measured equivalent width and associated random and systematic errors. The latter were determined by shifting the continuum by \( \pm \sigma / \sqrt{n} \) (where \( \sigma \) is the 1\( \sigma \) error spectrum, and \( n \) is the number of independent resolution elements over which the equivalent width integration was carried out), and recalculating the equivalent width. For the undetected transitions, I quote the 3\( \sigma \) limiting rest-frame equivalent width, using as a guide the profile of the weakest absorption line, Fe II \( \lambda 1608 \), which I detect at the 5\( \sigma \) confidence limit (see Table 5.1). Examples of metal absorption lines are reproduced in Figure 5.1.

As can be appreciated from inspection of Table 5.1 and Figure 5.1, the metal lines in this
5.3 Profile Fitting and Abundance Analysis

Figure 5.1: Selected absorption lines in the $z_{\text{abs}} = 2.3400972$ DLA towards J0035$-$0918. The data are shown with black histograms, while the red continuous lines are model fits to the line profiles. Top panel: Portion of the MagE spectrum of J0035$-$0918 encompassing the damped Ly$\alpha$ line, together with the theoretical Voigt profile (red line) for a neutral hydrogen column density $\log[N(\text{HI})/\text{cm}^{-2}] = 20.55 \pm 0.1$. The remaining panels display portions of the HIRES spectrum near metal lines of interest, together with model profiles generated with VPFIT as described in Section 5.3.1. This DLA consists of a single absorption component at $z_{\text{abs}} = 2.3400972$ with a small velocity dispersion, $b = \sqrt{2}\sigma = 2.4$ km s$^{-1}$. The weak absorption feature centred at $+26$ km s$^{-1}$ in the Si$\text{II}\lambda 1260.4221$ panel is probably Fe$\text{II}\lambda 1260.533$ absorption in the DLA, although its strength is below the 3$\sigma$ detection limit. The red wings of both C$\text{II}\lambda 1036$ and O$\text{I}\lambda 988$ are blended with a weak Ly$\alpha$ forest line indicated by a continuous blue line. In all plots the $y$-axis scale is residual intensity.
DLAs are very weak and narrow, with equivalent widths $W_0 < 55$ mÅ (the strongest line being C II $\lambda$1334), and with the absorption taking place in a single velocity component with FWHM < 10 km s$^{-1}$. The weakness of the absorption limits the detection to the intrinsically most abundant elements of the periodic table, C, N, O, Al, Si, and Fe; on the other hand, the wide wavelength coverage of the echelle spectra, which reach well into the far-UV, gives access to several transitions of differing $f$-values for most of these elements.

### 5.3.1 Column Densities

I begin the abundance analysis by measuring the column density of neutral hydrogen. To this end, I used the MagE spectrum of the QSO, the relevant portion of which is reproduced in the top panel of Figure 5.1, because at $z_{\text{abs}} = 2.3400972$ the damped Ly$\alpha$ line falls on a gap between two of the CCDs in the HIRES detector mosaic. Even at the coarser resolution of MagE (compared to HIRES), the broad damped Ly$\alpha$ line is fully resolved and no loss of accuracy results in the derivation of $N$(H I). Fitting a Voigt profile to the line, I deduced $\log[N$(H I)/cm$^{-2}] = 20.55 \pm 0.10$; the corresponding theoretical Voigt profile is overlaid on the MagE spectrum in the top panel of Figure 5.1.

In the next step, I determined the Doppler parameter of the absorbing gas, $b$ (km s$^{-1}$), and the column density of the metal ions, $N$(X) (cm$^{-2}$), by fitting the corresponding line profiles with VPFIT$^3$, which simultaneously fits multiple Voigt profiles to several atomic transitions, returning the values of $N$(X) and $b$, together with the associated errors, that minimise the $\chi^2$ between the data and the model. I tied the redshift and the Doppler parameter to be the same for all of the absorption lines listed in Table 5.1, which is justified if the neutrals and first ions are kinematically associated with the same gas (I relax this assumption later). With these constraints, VPFIT converged to a best-fitting model consisting of a single absorption component with redshift $z = 2.3400972 \pm 0.0000008$ and Doppler parameter $b = 2.36 \pm 0.08$ km s$^{-1}$. The corresponding column densities are listed in Table 5.2. Examples of the theoretical line profiles generated by VPFIT are shown superimposed on the data in Figure 5.1.

The weakest feature in the data is S II $\lambda$1259. When this line is included in the VPFIT fitting procedure (see bottom right panel of Figure 5.1), I derive a column density $\log[N$(S II)/cm$^{-2}] = 13.08 \pm 0.10$. However, since this absorption line is only significant at the $\sim 4.5\sigma$ level, I conservatively consider the above value to be an upper limit to the column density of S II.

With the usual assumption that the ions observed are the dominant stage of the corresponding elements in H I gas, so that corrections for unseen ion stages and/or the presence of ionised gas are negligible (I review this assumption in Section 5.3.2 below), it is straightforward to deduce the abundances of the elements concerned by simply dividing the column densities in Table 5.2 by $N$(H I). Comparison with the solar abundance scale of Asplund et al. (2009) then leads to the abundance pattern listed in Table 5.3 and illustrated graphically in Figure 5.2.

---

$^3$VPFIT is available from http://www.ast.cam.ac.uk/~rfc/vpfit.html
5.3 Profile Fitting and Abundance Analysis

Table 5.2: Ion column densities in the \( z = 2.3400972 \) DLA towards J0035−0918

<table>
<thead>
<tr>
<th>Ion</th>
<th>( \log N(X)/\text{cm}^{-2} )</th>
</tr>
</thead>
<tbody>
<tr>
<td>H I</td>
<td>20.55 ± 0.10</td>
</tr>
<tr>
<td>C II</td>
<td>15.47 ± 0.15</td>
</tr>
<tr>
<td>N I</td>
<td>13.51 ± 0.06</td>
</tr>
<tr>
<td>O I</td>
<td>14.96 ± 0.08</td>
</tr>
<tr>
<td>Al II</td>
<td>11.73 ± 0.05</td>
</tr>
<tr>
<td>Si II</td>
<td>13.41 ± 0.04</td>
</tr>
<tr>
<td>S II</td>
<td>≤ 13.08</td>
</tr>
<tr>
<td>Fe II</td>
<td>12.98 ± 0.07</td>
</tr>
</tbody>
</table>

Table 5.3: Element abundances in the \( z = 2.3400972 \) DLA towards J0035−0918

<table>
<thead>
<tr>
<th>Element</th>
<th>( \log \varepsilon(X)_{\text{DLA}} )</th>
<th>( \log \varepsilon(X)_{\odot} )</th>
<th>( [X/H]_{\text{DLA}} )</th>
</tr>
</thead>
<tbody>
<tr>
<td>C</td>
<td>6.92</td>
<td>8.43</td>
<td>−1.51</td>
</tr>
<tr>
<td>N</td>
<td>4.96</td>
<td>7.83</td>
<td>−2.87</td>
</tr>
<tr>
<td>O</td>
<td>6.41</td>
<td>8.69</td>
<td>−2.28</td>
</tr>
<tr>
<td>Al</td>
<td>3.18</td>
<td>6.44</td>
<td>−3.26</td>
</tr>
<tr>
<td>Si</td>
<td>4.86</td>
<td>7.51</td>
<td>−2.65</td>
</tr>
<tr>
<td>S</td>
<td>≤ 4.53</td>
<td>7.14</td>
<td>≤ −2.61</td>
</tr>
<tr>
<td>Fe</td>
<td>4.43</td>
<td>7.47</td>
<td>−3.04</td>
</tr>
</tbody>
</table>

\( a \log \varepsilon(X) = 12 + \log N(X)/N(\text{H}) \).

\( b \) Reproduced from Appendix A.

\( c [X/H]_{\text{DLA}} \equiv \log \varepsilon(X)_{\text{DLA}} - \log \varepsilon(X)_{\odot} \), with errors as listed in Table 2.

5.3.2 Ionisation Corrections

In DLAs, it is usually assumed that the metals within the absorbing H I gas reside in a single dominant ionisation stage, \( X_N \), so that the total abundance of an element is given by

\[
[X/H] = [X_N/H_I] + IC(X)
\]  

(5.1)

where the ionisation correction, IC(X), is typically negligible. In general, it is safe to assume IC(X) \( \simeq 0.0 \) for gas with high \( N(\text{H}_I) \), since the gas is self-shielded from ionising radiation (e.g. Vladilo et al., 2001). Of course, if the gas does not reside in a single dominant ionisation stage, or some amount of \( X_N \) is associated with H II gas, one may respectively under- or over-estimate the abundance of a given element (positive or negative IC(X) respectively).

To gauge the extent of such corrections, I used the CLOUDY photoionisation software (Ferland et al., 1998) to model the DLA as a slab of gas with uniform density in the range \( −3 < \log[n(\text{H})/\text{cm}^{-3}] < 3 \), exposed to the Haardt & Madau (2001) metagalactic ionising background and the cosmic microwave background, both at the redshift of the DLA. Adopting the solar abundance scale listed in Appendix A, I globally scaled the metals to \( \log Z_{\text{DLA}}/Z_\odot = −2.75 \) (an approximate average metallicity—see Figure 5.2). Once the column density of the DLA was reached, I stopped the simulations and output the resulting ion column densities. Using these

Figure 5.2: Element abundances in the \( z_{\text{abs}} = 2.3400972 \) DLA towards J0035−0918. The height of each box represents the uncertainty in each element abundance; the 3\( \sigma \) upper limit for S is indicated by the bar and arrow.
The results from these CLOUDY simulations are shown in Figure 5.3. The ionisation corrections appropriate to the DLA under investigation depend on the volume density of the gas, which can be inferred from the ratio of successive ion stages (see right panel of Figure 5.3). The only element for which I have this information is nitrogen and, even then, I can only derive a 3σ upper limit to the N II column density from the 3σ upper limit on the rest frame equivalent width of the undetected N II λ1084 line (see Table 5.1), leading to the upper limit log N(N II)/N(N I) ≤ −0.91. This affords a lower limit on the volume density of log [n(H)/cm⁻³] ≥ −1.0 (see right-hand panel of Figure 5.3). Referring now to the left-hand panel of Figure 5.3, it can then be seen that, when the gas density is greater than 0.1 cm⁻³, the ionisation corrections for the ions of interest here are indeed small, IC < ∼ 0.1 dex. These values are comparable to the uncertainties in the ion column densities (see Table 5.2) which justifies my assumption that [X/H] ∼ [X N/H I], and that ionisation corrections are not a serious concern.

It is worth noting, in passing, that even in the absence of several density-sensitive ion ratios, low values of n(H) are in general unlikely in these metal-poor DLAs with such simple velocity structure. The column density log[N(H I)/cm⁻²] = 20.55 implies a linear size l > 1 kpc along the line of sight if log[n(H)/cm⁻³] < −1. It seems unlikely that the structures giving rise to DLAs with properties similar to that considered here could have such large physical extent while maintaining very quiescent kinematics, with internal velocity dispersions of only a few km s⁻¹.
5.4 A masquerading carbon enhancement?

Returning to Figure 5.2, I note that with \([\text{O/H}] = -2.28\) and \([\text{Fe/H}] = -3.04\) the \(z = 2.3400972\) DLA in line to J0035–0918 is among the most metal-poor known (Pettini et al., 2008a; Pen-prase et al., 2010). However, the most striking feature of the abundance pattern in Figure 5.2 is the overabundance of carbon relative to all other elements. Thus, for example, \([\text{C/O}] = +0.77\), which implies an overabundance of carbon relative to oxygen by a factor of \(\sim 6\); similarly, \([\text{C/Fe}] = +1.53\), or \(\sim 35\) times solar! Given such extreme values, it is important to consider what factors, if any, may have resulted in a spurious overestimate of the carbon abundance. To this end, I performed several tests, which I now discuss.

5.4.1 Incorrect background subtraction?

First I consider the possibility that the background may have been oversubtracted in the proximity of the \(\text{C}\,\text{II}\) absorption lines; if the zero level had been incorrectly determined, this may lead one to overestimate the apparent optical depth of the absorption. While I cannot categorically rule out this possibility, as I do not have any independent measures of the zero level (such as saturated absorption lines with flat cores) in the immediate vicinity of the \(\text{C}\,\text{II}\) lines, I note the following: (i) there are a number of strongly saturated Ly\(\alpha\) absorption lines in the Ly\(\alpha\) forest (shortwards of \(\lambda_{\text{obs}} = 4158\,\text{Å}\)) and none of them show a systematic offset of their cores from zero residual intensity; (ii) the apparent optical depths of the two \(\text{C}\,\text{II}\) transitions covered by the HIRES spectrum, \(\lambda 1334\) and \(\lambda 1036\), are mutually consistent (see Figure 5.1); thus, if the background level were incorrect, it would have to have been oversubtracted by the same fractional amount of the QSO continuum at these two wavelengths, separated by \(995\,\text{Å}\) in the observed spectrum (at \(z_{\text{abs}} = 2.3400972\)). I also inspected the raw 2-D HIRES frames in the region of the two best observed \(\text{C}\,\text{II}\) and \(\text{O}\,\text{I}\) lines, \(\lambda 1334\) and \(\lambda 1302\) and found that they both fall close to the peak of the echelle blaze function, near the centre of the HIRES detector, where the data are of the highest S/N. I conclude that the recorded excess of C is unlikely to be an artifact of the data reduction process.

5.4.2 Profile fitting 1: gas kinematics

Next I looked critically at the profile fitting procedure. My \textsc{vpfit} modelling described in Section 5.3.1 assumed that the neutrals and singly ionised species arise from gas with the same Doppler parameter. I relaxed this constraint by fitting separately the absorption lines of the first ions (two \(\text{C}\,\text{II}\), five \(\text{Si}\,\text{II}\) lines and one line each of \(\text{Al}\,\text{II}\) and \(\text{Fe}\,\text{II}\)) and the neutrals (five \(\text{O}\,\text{I}\) and two \(\text{N}\,\text{I}\) lines). In this case, \textsc{vpfit} converged to best fitting Doppler parameters of \(b = 2.30 \pm 0.09\) for the first ions and \(b = 2.6 \pm 0.2\) for the neutrals. With these parameters the \(\text{C}\,\text{II}\) column density is higher by 0.1 dex than the value listed in Table 5.2 (that is, the value obtained by tying \(b\) to be the same for all species), while \(N(\text{O}\,\text{I})\) is lower by 0.1 dex. The column
densities of other elements remain essentially unchanged. In other words, kinematically decoupling neutrals and first ions has the net effect of further increasing the \([C/O]\) overabundance relative to the values in Table 5.3 and Figure 5.2. Note also that the above values of \(b\) are still consistent within the errors with the best-fitting \(b = 2.36 \pm 0.08\) deduced in Section 5.3.1 by tying all the absorption lines together. Finally, fitting only the \(\text{C}\,\text{II}\) lines, without reference to any other transition, yielded \(\log[\mathcal{N}(\text{C}\,\text{II})/\text{cm}^{-2}] = 15.42 \pm 0.59\), which is only 0.05 dex lower than the value listed in Table 5.2, albeit with a larger error (but with much the same redshift and Doppler parameter: \(z_{\text{abs}} = 2.340096 \pm 0.000002\) and \(b = 2.4 \pm 0.3 \text{ km s}^{-1}\)).

5.4.3 Profile fitting 2: instrumental resolution

In order to compare theoretical and observed absorption line profiles, VPFIT requires knowledge of the instrumental broadening function, normally assumed to be a Gaussian with FWHM corresponding to the spectral resolution. The nominal resolution of spectra recorded through the 1.148 arcsec wide entrance slit of HIRES employed in my observations is FWHM = 8.1 km s\(^{-1}\) (see http://www2.keck.hawaii.edu/inst/hires/), but this applies to a uniformly illuminated slit. Since the seeing was consistently better than 1.148\(\prime\) during the observations of J0035\,−\,0918, it is likely that the actual resolution of the data is somewhat better than the nominal value. The narrowest features in the spectrum of J0035\,−\,0918 are the metal absorption lines of the \(z_{\text{abs}} = 2.3400972\) DLA themselves. I therefore estimated the true spectral resolution—measured by the Doppler parameter \(b_{\text{instr}} \equiv 0.6006\) FWHM—by varying \(b_{\text{instr}}\) in small steps from 5.0 to 3.7 km s\(^{-1}\) (i.e. between FWHM = 8.3 and 6.2 km s\(^{-1}\)), fitting all of the absorption lines as described in Section 5.3.1, and minimising the \(\chi^2\) between the model and observed profiles as a function of \(b_{\text{instr}}\). This procedure gave a well-defined \(b_{\text{instr,min}} = 4.4\) km s\(^{-1}\), corresponding to FWHM = 7.3 km s\(^{-1}\), or \(R = 41000\), which is the value used in all the profile fitting described in Sections 5.3.1 and 5.4.2.

It is obviously important to test the sensitivity of these results to the value of \(b_{\text{instr}}\) adopted. Within the range of values tested, \(b_{\text{instr}} = 5.0–3.7\) km s\(^{-1}\), the column density of \(\text{C}\,\text{II}\) and the corresponding element ratios changed by no more than \(\pm 0.1\) dex compared to the values in Table 5.2 which refer to the case \(b_{\text{instr,min}} = 4.4\) km s\(^{-1}\).

5.4.4 Profile Fitting 3: thermal broadening of the line profiles

When deriving the cloud parameters detailed in Section 5.3.1, I assumed the gas that gives rise to the DLA absorption has a temperature of 100 K, so that the broadening of the line profile is dominated by the turbulent motions of the gas (i.e. \(b_{\text{th}}^2 \ll b_{\text{turb}}^2\); cf. Eq. 2.5 and 2.6). If, however, the gas temperature is somewhat larger than 100 K, so that \(b_{\text{th}}^2 \sim b_{\text{turb}}^2\), then the total Doppler parameter will be narrower for heavier elements, acting to reduce the \([\text{C}/\text{Fe}]\) ratio.

Unfortunately, the effects of thermal broadening would produce very small deviations to
Figure 5.4: The \( \chi^2 \) output by VPFIT is plotted for a range of plausible gas temperatures for the DLA (top left panel). Note that there are a total of 151 degrees of freedom. VPFIT favours a solution corresponding to the lowest permitted gas temperature, although the \( \chi^2 \) is essentially flat below \( \sim 6000 \) K. When \( \chi^2 \approx 163.0 \) (corresponding to the 2\( \sigma \) limiting contour for the \( \chi^2 \); see text for further details), I infer a 2\( \sigma \) upper limit of \( T_{\text{gas}} \) \((2\sigma) < \sim 8500 \) K, represented by the vertical dashed line. In the remaining panels, I show how the variations in the gas temperature affect the derived C\( \text{II} \) column density (top right panel), [C/Fe] abundance (bottom left panel), and [C/O] abundance (bottom right panel).

the already unresolved line profiles; the gas temperature cannot be measured directly from the present data. Instead, I used VPFIT to determine an upper limit for the DLA’s gas temperature, and thus a lower limit on the [C/Fe] ratio, as follows. Given that the largest variations will be exhibited by the relatively low mass ions, I only used as input into VPFIT the measured line profiles of C\( \text{II} \) and O\( \text{I} \), having fixed the redshift to the value derived in Section 5.3.1 (i.e. \( z_{\text{abs}} = 2.3400972 \)). By fixing the gas temperature, I allowed VPFIT to derive the best-fitting turbulent Doppler parameter, and column density for C\( \text{II} \) and O\( \text{I} \). This process was then repeated for different values of the gas temperature. Thus, there are four parameters that were adjusted to determine the minimum \( \chi^2 \); \( N(\text{C\( \text{II} \)),} \) \( N(\text{O\( \text{I} \)),} \) \( b_{\text{turb}}, \) and \( T_{\text{gas}} \). Referring to the top left panel of Figure 5.4, it can be seen that VPFIT favours the minimum allowable gas temperature, corresponding to a \( \chi^2_{\text{min}}/\text{dof} \approx 153.5/151 \). One can now determine the 2\( \sigma \) limiting contour for the \( \chi^2 \) from a standard \( \chi^2 \) table (see e.g. Wall & Jenkins 2003); for the four adjustable parameters used here, \( \chi^2(0.05) = 9.49 \). Thus, the limiting value is \( \chi^2_{\text{lim}} = \chi^2_{\text{min}} + \chi^2(0.05) \approx 163.0 \) (see e.g. Lampton, Margon, & Bowyer 1976), corresponding to a 2\( \sigma \) upper limit to the
gas temperature of \( T_{\text{gas}}(2\sigma) \approx 8500 \text{ K} \). In the remaining panels of Figure 5.4, I illustrate how the other fitted parameters vary with \( T_{\text{gas}} \). For the derived 2\( \sigma \) upper limit on the gas temperature, these plots suggest that \( \log N(\text{C II})/\text{cm}^{-2} \geq 14.4 \), \([\text{C/Fe}] \geq +0.46\), and \([\text{C/O}] \geq -0.03\).

Even in the event of such extreme conditions, the \([\text{C/Fe}]\) ratio is still a factor of two higher than any other DLA known (Pettini et al., 2008a). In all likelihood, the DLA contains both a warm (a few \( \times 10^3 \text{ K} \)) and a cold (a few \( \times 10^2 \text{ K} \)) neutral medium. A tighter constraint on the gas temperature could in principle be afforded in future from higher resolution spectra with access to the neutral deuterium lines. Nevertheless, on the basis of the tests carried out in this and the two preceding subsections, I conclude that the overabundance of carbon is not an artifact of the profile fitting analysis of the absorption lines.

### 5.4.5 Monte Carlo simulations

The \( z = 2.3400972 \) DLA in J0035−0918 is unique so far in showing such a marked overabundance of carbon. More typically, DLAs with \([\text{O/H}] < -2\) have \([\text{C/O}] \approx -0.2\) and at most \([\text{C/O}] \approx 0.0\) [i.e. \((\text{C/O}) = (\text{C/O})_\odot\); Pettini et al. 2008a, Chapter 4]. It is therefore worthwhile considering to what extent a solar C/O ratio is excluded by the data. I illustrate this test in Figure 5.5, where I compare the line profiles for the two C II absorption lines generated with: (upper panels) the best-fitting model returned by VPFIT, and (lower panels) a model in which \( N(\text{C II}) \) has artificially been fixed at the column density corresponding to \([\text{C/O}] = 0.0\) (that is \( \log [N(\text{C II})/\text{cm}^{-2}] = 14.70 \) as opposed to 15.47 as in Table 5.2). While the former results in a \( \chi^2/\text{dof} = 13.9/22 \), the latter fits are worse, with \( \chi^2/\text{dof} = 86.6/22 \). The higher value of \( \chi^2 \) is reflected by the higher residuals, both in the core and the wings of the C II \( \lambda 1334 \) line in particular, as can be appreciated from close examination of Figure 5.5. The noisier C II \( \lambda 1036 \) line is less instructive in this context, partly because of blending with Ly\( \alpha \) forest lines.

It may be surprising to some to see what subtle changes in the line profiles result from changing \( N(\text{C II}) \) by a factor of \( \sim 6 \). The reason is that the two C II absorption lines are close to saturation and, as they approach the flat part of the curve of growth, their equivalent widths only increase slowly with rising column density. In the present case, the curve of growth for the metal lines in the \( z_{\text{abs}} = 2.3400972 \) DLA towards J0035−0918 (reproduced in Figure 5.6) is well-constrained by the relative strengths of six O I and five Si II transitions of widely differing \( f \)-values (see Table 5.1). Since the column densities of both O I and Fe II are fixed by optically thin transitions (i.e. independent of the Doppler parameter), the ratios \([\text{C/O}]\) and \([\text{C/Fe}]\) depend only on the C II column density. In the right panel of Figure 5.6, I compare the measured equivalent width of C II \( \lambda 1334 \) (indicated by the blue hatched region) with the values expected if \([\text{C/O}] = 0.0\) (red), \([\text{C/Fe}] = +1.0\) (blue), and \([\text{C/Fe}] = +1.53\) (green). Only in this last case (\([\text{C/Fe}] = +1.53\)) do I recover the measured \( W_0(\text{C II} \lambda 1334) \), whereas the strength of the line is underpredicted in the other two cases illustrated. The partially blended C II \( \lambda 1036 \) line is not instructive in this equivalent width test which, by its nature, is less sensitive than the pixel-by-
5.4 A masquerading carbon enhancement?

Figure 5.5: Comparison between observed C\textsc{ii} line profiles (black histograms) and theoretical profiles computed with \textsc{vpfit} (red continuous lines). Below each plot I also show the difference between computed and observed residual intensities (red lines) compared with the 1\sigma error spectrum of the data (grey area). (a): best fitting model as described in Section 5.3.1 with the log\([N(C\textsc{ii})/cm^{-2}] = 15.47\) as in Table 5.2. (b): best fitting model obtained by forcing log\([N(C\textsc{ii})/cm^{-2}] = 14.70\) so that [C/O] = 0.0 (see text for further details). This model provides a poorer fit to both C\textsc{ii} absorption lines, as demonstrated by the mismatch between the model subtracted from the data and the 1\sigma error spectrum (bottom panels).

Figures 5.5 and 5.6 offer a clearer view of the carbon-enhanced nature of this DLA, while at the same time highlighting the subtle differences in the absorption line profiles when lines are on the flat part of the curve of growth. Thus, it is reasonable to question whether the differences between the line profiles computed with [C/O] = 0.0 and the observed profiles, most evident for C\textsc{ii} \(\lambda 1334\) in the bottom right-hand set of panels in Figure 5.5, may be simply due to statistical fluctuations. I tested this possibility with a Monte Carlo-type approach. I synthesised 100 pairs of C\textsc{ii} \(\lambda 1036, \lambda 1334\) line profiles, all corresponding to [C/O] = 0.0 and with the best fitting Doppler parameter \((b = 2.36 \pm 0.08)\). I perturbed these line profiles with Gaussian distributed errors from the observed 1\sigma error spectrum, and then used \textsc{vpfit} to fit each of these 100 pairs...
of C II lines (together with the observed line profiles of the other ions) and deduce 100 new values of \(N(C\text{II})\). Using the simultaneously derived O I column density (which is essentially unchanged since it is constrained by weaker transitions), I calculated the [C/O] ratio for all of my realisations to determine how often one would expect to observe [C/O] \(\gtrsim 0.77\) if the true [C/O] ratio were solar. The results of this Monte Carlo approach are shown in Figure 5.7. In none of the simulations was a value of [C/O] as high as +0.77 recovered. The distribution of recovered values is symmetric about the input value [C/O] = 0.0 with a dispersion \(\sigma = 0.087\). Thus, if the [C/O] ratio in the \(z = 2.3400972\) DLA were indeed solar (which in itself is higher or as high as observed in any other DLA up to now), the data would constitute a \(\sim 9\sigma\) deviation, which is very unlikely indeed.

In conclusion, none of the tests I have performed can explain away the C overabundance I have uncovered. The most straightforward interpretation of the data is that I have identified a damped Ly\(\alpha\) system with a marked overabundance of C, comparable to that so far found only in carbon-enhanced metal-poor stars of the Milky Way.

### 5.5 A Carbon-Enhanced metal-poor DLA

As already mentioned, this is the first example of a DLA meeting the conventional definition of CEMP stars, [C/Fe] \(> +1.0\). While it is the only such case in the survey presented in Chapter 4, which includes a dozen DLAs with [Fe/H] \(< -2\), it is nevertheless very hard to draw conclusions regarding the frequency of CEMP DLAs. The reason stems from the difficulties associated with measuring the C II column density; the C II \(\lambda 1036\) and \(\lambda 1334\) absorption lines are almost
5.5 A Carbon-Enhanced metal-poor DLA

Figure 5.7: Histogram of the values of $[\text{C/O}]$ recovered from 100 Monte Carlo realisations with input $[\text{C/O}]=0.0$ in which the synthetic C II $\lambda 1334, \lambda 1036$ line profiles were perturbed with a random realisation of the error spectrum. The overlaid Gaussian fit to the results has a dispersion $\sigma = 0.087$, which implies that the measured value $[\text{C/O}] = +0.77$ would be a $\sim 9\sigma$ random fluctuation if the $[\text{C/O}]$ ratio were solar in the $z = 2.3400972$ DLA.

always saturated in damped Ly$\alpha$ absorption systems, even at low metallicities. CEMP DLAs may thus be easily overlooked.

The discovery of a gaseous, high redshift, counterpart to CEMP stars of the Galactic halo is a breakthrough with potentially very important consequences. As mentioned in Section 5.1, interpretations still differ as to the origin of the carbon enhancement in very metal-poor stars, but the favoured explanation involves mass transfer from an AGB companion, at least for the members of the CEMP class which also exhibit enhancements in s-process elements. The situation is less clear-cut for the CEMP stars with no such enhancement, the CEMP-no stars. On the one hand, there may be a continuous link between this class and the CEMP-s class (Masseron et al., 2010), so that in some CEMP-no stars the carbon enhancement may also be due to mass transfer from an unseen companion. On the other hand, the binary transfer scenario cannot readily explain why the fraction of all metal-poor stars that exhibit a carbon-enhancement increases with decreasing metallicity (Beers & Christlieb, 2005). In fact, the only physical model so far put forward that can explain this increased fraction requires a high carbon yield from metal-free stars to efficiently cool low metallicity gas and drive the transition from Pop III to Pop II star formation (Frebel, Johnson, & Bromm, 2007).

While CEMP-no stars in the Galactic halo may have more than one origin, the same ambiguities do not apply to the carbon enhancement in the DLA presented here, where an entire interstellar cloud is overabundant in carbon (below I place limits to the total mass of carbon involved). Rather, what is presumably seen in this DLA is the initial chemical composition of the gas from which CEMP stars would subsequently form. In this case, the pattern of element abundances revealed gives strong clues to the nature of the earlier generation of stars (and supernovae) which seeded the DLA cloud with its metals.

I explore these ideas further by comparing the abundance pattern of the DLA with those of halo stars of similar metallicity and with calculations of element yields from metal-free and metal-poor stars. In this endeavour, I am limited by the small number of elements whose abun-
dances I have been able to measure, but this is an inevitable consequence of the low metallicity of this DLA. While it would be instructive to know the abundances of other Fe-peak elements, such as Cr and Zn for example, the absorption lines of these species would require S/N $> 500$ for $5\sigma$ detections! Realistically, only Mg and S are within reach of dedicated future observations, and there is certainly no hope of measuring directly the abundances of neutron capture elements in the DLA.

### 5.5.1 Comparison with Stellar Abundances

In order to compare the relative abundances of the elements measured in the DLA with the corresponding values in Galactic halo stars, I median-combined the abundances of all CEMP stars with Fe/H within a factor of two of the DLA ($-3.34 \leq [\text{Fe}/\text{H}] \leq -2.74$) from the recent compilation of Frebel et al. (2010). I then calculated the dispersion in each abundance using the Interactive Data Language routine `ROBUST_SIGMA` which determines the median absolute deviation (unaffected by outliers) of a set of measurements, and then appropriately weights the data to provide a robust estimate of the sample dispersion (Hoaglin, Mosteller, & Tukey 1983, cf. Chapter 3).

Figure 5.8 illustrates three comparisons. In the top panel, the DLA abundance pattern (solid black boxes) is shown together with that of the ‘average’ population of all CEMP stars (open magenta boxes), while in the middle panel the comparison is restricted to CEMP-no stars. Clearly CEMP stars, of both flavours, are a heterogeneous population, exhibiting wide ranges in the ratios of the elements considered here. On the basis of the limited evidence available, it appears that the C/Fe ratio in the DLA is a better match to the values seen in the subset of CEMP-no stars than to the more extreme values encountered in some CEMP stars. Other element ratios are less straightforward to interpret.

Nitrogen, for example, is evidently less abundant in the DLA compared to CEMP stars by $\sim 2$ orders of magnitude! However, it is difficult to draw firm conclusions from this discrepancy, because the abundance of N is notoriously difficult to measure in metal-poor stars where it relies on the analysis of molecular bands of NH or CN, each with its own disadvantages (see e.g. Asplund 2005). The largest uncertainty arises from the assumed 1D (rather than 3D) model atmospheres; correcting for 3D effects can change the deduced values of N/H by $\sim -0.5$ dex (Note that similar corrections also apply to C/H when derived from molecular bands of CH or C$_2$; Collet, Asplund, & Trampedach 2007). Furthermore, the photospheric N/H that is measured today is unlikely to reflect the N abundance of the gas cloud that gave birth to the star; during the star’s evolution it will become self-polluted by its own nucleosynthesis (e.g. through CN cycling or rotational mixing), which is likely emphasised by the high seed C abundance. Thus, a comparison between stellar and DLA nitrogen abundances is not very instructive.

---

$^4$ Assuming $[\text{Cr}/\text{Fe}] = [\text{Zn}/\text{Fe}] = 0.0$.

$^5$ Available from http://idlastro.gsfc.nasa.gov/homepage.html
Figure 5.8: Comparison of element abundances in the $z_{\text{abs}} = 2.3400972$ DLA (filled black boxes) and in Galactic halo stars with an Fe abundance within a factor of two of the DLA (open magenta boxes). The numbers below the element labels indicate the number of stars that contributed to the determination of the ‘typical’ stellar abundances, and the heights of the magenta boxes reflect the dispersion of each set of measurements. Top Panel: comparison with all CEMP stars that have $-3.34 \leq [\text{Fe}/\text{H}] \leq -2.74$. Middle Panel: comparison with CEMP-no stars that have $-3.34 \leq [\text{Fe}/\text{H}] \leq -2.74$. For this case, the oxygen abundance of a single CEMP-no star is shown by the open circle. Lower Panel: comparing the DLA abundance pattern with the stellar abundance patterns of HE 0143−0441 (red symbols; a CEMP-s star with $[\text{Fe}/\text{H}] = -2.21$, $[\text{Ba/Fe}] = +0.62$ from Cohen et al. 2004) and BD+44°493 (green symbols; a CEMP-no star with $[\text{Fe}/\text{H}] = -3.73$, $[\text{Ba/Fe}] = -0.55$ from Ito et al. 2009). Note that in this last panel, I have plotted $[\text{X}/\text{Fe}]$ as opposed to $[\text{X}/\text{H}]$. In all panels the dashed line represents the solar abundance.
Comparing the oxygen abundances in CEMP stars and the DLA is similarly problematic. The statistics are very poor (see Figure 5.8) and it is well known that different spectral features used to deduce O/H in the metal-poor regime give discordant answers (García Pérez et al., 2006). Significant departures from local thermodynamic equilibrium (LTE) become important when measuring the O abundance from the infrared triplet at 777 nm in the metal-poor regime ([Fe/H] < −2.5), leading to non-LTE corrections of the order of −0.5 to −1.0 dex (Fabbian et al., 2009b). Similarly, the oxygen abundance determined from the UV OH lines is subject to large 3D corrections of up to −0.9 dex when [Fe/H] ≤ −3.0 (Asplund, 2005). Clearly, it will be important to reexamine this issue once samples of CEMP stars have grown to include more cases where [O/Fe] has been measured with confidence.

Finally, Figure 5.8 demonstrates that there is a reasonably good match between CEMP stars and the DLA in the elements Al and Si, particularly for the CEMP-no stars. In this context, I point out that significant (positive) non-LTE corrections to Al/H may apply for stars in this metallicity regime (see e.g. Asplund 2005). To address this issue, I first identified those CEMP stars in the Frebel et al. (2010) compilation whose Al abundances were derived assuming LTE. I then applied non-LTE corrections to Al based on the Andrievsky et al. (2008) estimates, which amounted to a typical correction of the order of +0.6. The Al abundances for the ‘typical’ CEMP and CEMP-no star presented in Figure 5.8 should thus correspond to the (approximate) non-LTE values.

Finally, in the lower panel of Figure 5.8 I compare the element ratios (relative to Fe) in the DLA with the values measured for two CEMP stars selected from the compilation by Frebel et al. (2010) because they most closely match the relative abundances in the DLA. It is certainly plausible that stars forming out of this DLA gas would share many chemical similarities – at least for the elements considered here – with these two stars of the Milky Way Galactic halo.

### 5.5.2 Comparison with Model Yields for Metal-free Stars

With a metallicity [Fe/H] of only 1/1000 solar, it is conceivable that the chemical composition of the DLA gas reflects the element yields of only a few prior generations of stars. It is thus of interest to compare the abundance pattern in Figure 5.2 with calculations of nucleosynthesis by metal-free, or low metallicity, stars. Calculations of yields from metal-free stars (commonly referred to as Pop III stars), aimed in particular at interpreting the element ratios seen in very metal-poor stars, have focused on models which include ‘mixing and fallback’ (e.g. Umeda & Nomoto, 2002, 2003; Heger & Woosley, 2010). These scenarios involve core-collapse supernovae where the elements synthesised in the inner regions of the star are mixed by some process (several possibilities have been put forward); a fraction of the mixed material subsequently falls back onto the central remnant while the rest is ejected into interstellar space.

Such models are illustrative, rather than predictive, since the parameters describing the boundaries of the mixing region and the fraction of the mixed material which is ejected cannot
be derived from first principles, and are instead parameterised and suitably adjusted to fit the observed stellar abundances (e.g. Tominaga et al. 2007). A string of recent models that employ a more physically motivated prescription of fallback (Heger & Woosley, 2010), have also investigated the effects of mixing due to the Rayleigh-Taylor instability (Joggerst, Woosley, & Heger, 2009) together with rotationally induced mixing in 2D (Joggerst et al., 2010a) and 3D (Joggerst, Almgren, & Woosley, 2010b). By and large, a general feature of these models is that an abundance pattern similar to that found here in the DLA, including the carbon enhancement and the marked odd-even effect (see Figure 5.2), can be reproduced with low energy explosions, such as those giving rise to the faint SN branch (see Figure 1 of Tominaga et al. 2007), and a moderate degree of mixing and fallback. For example, in the mixing-fallback model considered by Tominaga (2010), enrichment with the elements ejected by a single $25 M_\odot$ star of zero metallicity, exploding as a core-collapse supernova with an energy $E_{SN} \sim 10^{51}$ ergs (represented by the orange stars in Figure 5.9), provides a remarkably good fit to the relative element abundances measured in the DLA (see also Kobayashi, Tominaga, & Nomoto 2011).

It remains to be established whether a zero initial metallicity (i.e. Pop III) is actually required to match the observations, or whether a similar abundance pattern could also be reproduced by models with low, but non-zero, initial metallicity. An alternative to the mixing and fallback model has been proposed by the Geneva group whose work places more emphasis on the effects of rotation to provide the mixing and the trigger for mass loss in very metal-poor stars (e.g. Meynet, Ekström, & Maeder, 2006; Hirschi, 2007). Whilst these models are physically well-motivated, the rather high nitrogen yield predicted by some models contrasts with the relatively low [N/C] and [N/O] observed in this DLA. This may in turn place interesting constraints on the rotation velocities of the metal-free stars that may have seeded the DLA with its metals. Only recently has rotation been included in zero-metallicity progenitors under the mixing-fallback scenario (Joggerst et al., 2010a). It should be possible to assess better the effects of rotation on the yield from zero-metallicity progenitors as these stellar models improve and as the samples of CEMP DLAs grow.

Figure 5.9: A comparison of element abundances in the $z_{abs} = 2.3400972$ DLA (filled black boxes) with the model yield calculations from a $25 M_\odot$ Population III star that explodes as a CCSN with $10^{51}$ erg of energy (orange stars).
5.5.3 How Many Supernovae?

It is truly intriguing that an entire cloud of neutral gas is so highly enhanced in carbon and shows such a clear-cut odd-even effect, if both are indeed characteristic of Pop III supernova yields. In this context, it is of interest to estimate the mass of carbon involved. With the assumption of spherical symmetry, the mass of singly ionised carbon in the DLA is given by:

\[
M(C^{\text{II}}) = 12M(\text{H}) \frac{N(C^{\text{II}})}{N(\text{H})} \tag{5.2}
\]

\[
M(C^{\text{II}}) = 2\pi m_{\text{H}} \frac{N(\text{H})^2 N(C^{\text{II}})}{n(\text{H})^2} \tag{5.3}
\]

\[
M(C^{\text{II}}) \leq 200 \left( \frac{n(\text{H})}{0.1 \text{ cm}^{-3}} \right)^{-2} M_{\odot} \tag{5.4}
\]

where \(m_{\text{H}}\) is the mass of a hydrogen atom and \(n(\text{H})\) is the volume density (cm\(^{-3}\)). The corresponding mass of neutral gas in the DLA is:

\[
M_{\text{DLA}} = 1.3 m_{\text{H}} \frac{4\pi N(\text{H})^3}{3 \times 8 n(\text{H})^2} \tag{5.5}
\]

\[
M_{\text{DLA}} \leq 2.5 \times 10^6 \left( \frac{n(\text{H})}{0.1 \text{ cm}^{-3}} \right)^{-2} M_{\odot} \tag{5.6}
\]

Since C is mostly singly-ionised in DLAs, Eq. (5) implies an upper limit of \(200M_{\odot}\) to the total mass of \(^{12}\text{C}\) in the DLA. However, the inverse square dependence of \(M(C^{\text{II}})\) on the gas density allows substantially lower values of \(M(C^{\text{II}})\). In Section 5.3.2 it was shown that a lower limit of \(n(\text{H}) \geq 0.1 \text{ cm}^{-3}\) was based on the non-detection of N\(^{\text{II}}\) absorption, and that higher densities are likely given the very low velocity dispersion of the gas. If, for example, \(n(\text{H}) \approx 1 \text{ cm}^{-3}\), which would imply a linear extent of the DLA of 100 pc along the line of sight, then the implied mass of \(^{12}\text{C}\) would be reduced to only \(2M_{\odot}\). For comparison, the total mass of \(^{12}\text{C}\) ejected by the single 25\(M_{\odot}\) Pop III star in the model by Tominaga (2010) is \(\approx 0.2M_{\odot}\). The models by Heger & Woosley (2010) anticipate a similar ejected mass of \(^{12}\text{C}\) (\(\approx 0.3M_{\odot}\)) for a comparable explosion energy and progenitor mass. Thus, we may indeed be seeing the elements synthesised by only a few supernovae in the chemical enrichment of the DLA considered here.

In conclusion, I speculate that the \(z_{\text{abs}} = 2.3400972\) DLA in front of the QSO J0035–0918 may well be the much sought ‘missing link’ between the first, zero-metallicity, stars and the most metal-poor stars in the halo of our Galaxy of similar metallicity to the DLA under investigation. Its low metallicity of 1/1000 solar in Fe, coupled with an overabundance of C and a marked odd-even effect in the relative abundances of the few elements that could be measured, are all consistent with the yields produced by models of Pop III stars which explode as core-collapse supernovae of relatively low energy. The mass of newly synthesised elements may be that produced by only a few such supernovae, depending on the unknown volume density of the gas in the DLA. In this scenario, the gas seen as a damped Ly\(\alpha\) system at high redshift may be
the material from which a subsequent generation of stars formed, with a chemical composition similar to that seen in CEMP-no stars of the Galactic halo.

### 5.6 Summary and Conclusions

In the course of my programme to study the most metal-poor DLAs and, in particular, to measure the abundances of the CNO group of elements, I have uncovered a DLA, at $z = 2.3400972$ towards the $z_{\text{em}} = 2.42$ SDSS QSO J0035–0918, whose chemical composition is consistent with that produced from exploding Population III stars. From the analysis of medium and high resolution spectra of this DLA I draw the following conclusions.

(i) The metal absorption lines associated with the DLA are formed in very quiescent gas, consisting of a single absorption component with a small Doppler parameter $b = 2.4$ km s$^{-1}$.

(ii) The metallicity of the DLA, as measured from Fe, is very low: $[\text{Fe/H}] = -3.04 \pm 0.17$, or $\sim 1/1000$ solar.

(iii) The DLA exhibits an enhancement of the abundance of carbon relative to all other metals covered by the data: N, O, Al, Si, S and Fe. I measure $[\text{C/Fe}] = +1.53$, a factor of $\sim 20$ greater than observed in any other DLA up to now. Adopting the defining criterion for carbon-enhanced metal-poor stars in the Galactic halo, $[\text{C/Fe}] > +1.0$, this is the first example of an analogous carbon-enhanced DLA. I also deduce $[\text{C/O}] = +0.77$, whereas in all other DLAs with $[\text{Fe/H}] < -2$ studied up to now $[\text{C/O}] \lesssim 0.0$. One remaining concern is whether or not thermal broadening plays a significant role in broadening the line profiles. Under the assumption that thermal broadening dominates over turbulent broadening, I have inferred a strict lower limit on the C/Fe ratio of $[\text{C/Fe}] \geq +0.46$, which is still at least a factor of 2 greater than any other DLA that has been observed to date.

(iv) The DLA also exhibits a clear odd-even effect, which implies a low neutron excess, and hence presumably low abundances of neutron-capture elements. When its chemical composition is compared with that of Galactic carbon-enhanced metal-poor stars which do not exhibit an excess of neutron-capture elements, a good match is found for some element ratios (C/Fe, Al/Fe and Si/Fe). N and O are significantly less abundant, compared to Fe, in the DLA than in most CEMP stars, but it is difficult to draw definite conclusions for these two elements which are notoriously difficult to measure in very metal-poor stars.

(v) The abundance pattern I observe for this DLA is consistent with enrichment from a population of $\sim 20–40 M_\odot$ metal-free, or Population III, stars that ended their lives as core-collapse supernovae with modest explosion energies. I estimate the total mass of $^{12}\text{C}$ within the DLA to be $\leq 200 M_\odot$. This upper limit could be constrained further with higher signal-to-noise ratio spectra that would permit a measure of the column density of N II and other successive ion stages whose ratios are density dependent. The steep dependence of $M(^{12}\text{C})$ on the gas density allows the possibility that we may be seeing the chemical enrichment produced by only a few
prior supernovae.

**vi** I speculate that the gas in this DLA may be the ‘missing link’ between the yields of Population III stars, and their later incorporation in the CEMP-no class of carbon-enhanced metal-poor stars. I note, however, that the carbon-enhancement in some CEMP-no stars could also be produced by other means. Long term radial velocity monitoring of CEMP-no stars will confirm or deny their association with a now extinct binary companion.

The results presented here emphasise the importance of further observations of DLAs with $[\text{Fe/H}] \leq -3$ to complement the work being carried out on the most metal-poor stars. Such DLAs may be the most suitable environments for measuring the true yields from zero- or low-metallicity stars, free from the complications of stellar abundance measurements and the possibility that these stars are polluted by an unseen binary companion, or self-polluted by their own nucleosynthesis. The extremely metal-poor regime for DLAs is yet to be thoroughly explored with high-resolution and high signal-to-noise spectra—who knows what may lurk there!
Finding the First Metals

The gas that gives rise to damped Lyα systems probes a range of astrophysical environments; from the large, extended neutral gas discs of weakly star-forming galaxies, to their smaller constituent clouds of gas within. The haloes of these galaxies also contain low mass gas-rich satellites that can give rise to DLAs (see e.g. Pontzen et al. 2008), which were fed to the galaxy through a filament that, naturally, also contains such systems. The underlying goal of this thesis was to explore the properties of the most metal-poor DLAs, which likely arise in the halo of a galaxy or within some cosmic filamentary structure. The cold kinematics of such DLAs in conjunction with their low metal-content imply that these systems are typically of small physical extent, and form stars at either low rates, or perhaps not at all (see Chapter 3). This brings into question the origin of their metals; what stellar sources enriched the most metal-poor DLAs? Did the metals come from a wind that was driven by the Population II stars in the host galaxy, or were the metals present in the IGM when the DLA first condensed? Could these DLAs solely contain the metals from metal-free Population III stars? If this is the case, were the most metal-poor DLAs enriched with the metals from an entire generation of stars that globally enriched the Universe at early times (Madau, Ferrara, & Rees, 2001) or by just a handful of sources locally at much later times (Bland-Hawthorn et al., 2011)?

Several recent studies suggest that the most metal-poor DLAs do indeed contain the first synthesised metals in the Universe, or in the very least, they still retain the nucleosynthetic fingerprints of metal-free stars (Erni et al., 2006; Pettini et al., 2008b). Indeed, the primary conclusion from the work presented in this thesis is that: “The most metal-poor DLAs act as an intermediary between the enrichment episodes of the first generation of stars, and the incorpo-
ration of these metals into the second generation of stars, like those in the halo of our Galaxy” (see Chapter 5). Unfortunately, we only have access to the most abundant metals in these near-pristine systems, and we must therefore rely on these few metals as indicators to the enrichment source. Quite often, it is difficult to discern between Population II and Population III enrichment with such few indicators (see Chapter 4), unless a system is uncovered (such as the one discussed in Chapter 5) where the elemental abundance pattern cannot possibly be reproduced by Population II nucleosynthesis calculations and yet is naturally explained by Population III star models. In this concluding Chapter, I outline the future work that might resolve some of the questions and difficulties raised above.

6.1 What fraction of DLAs are carbon-enhanced?

The origin of the carbon-enhanced signature that is observed at low metallicity is typically attributed to the Pop III stars that exploded with a relatively low explosion energy. This results in a large fraction of the $^{56}$Ni falling back onto the central remnant (see e.g. Umeda & Nomoto 2003), whilst virtually all of the carbon is ejected, since it is synthesised in the He-burning region in the outer stellar layers. Essentially, the degree to which a system is carbon-enhanced depends primarily on the final kinetic energy of the explosion at infinity (i.e. the energy of the ejected material). Thus, the fraction of carbon-enhanced DLAs may provide a loose constraint on the typical explosion energy of metal-free stars.

To demonstrate this point I used the STARFIT software provided by Heger & Woosley (2010), together with their extensive yield calculations for Pop III stars that end their lives as CCSNe, to extract all models that give rise to a carbon-enhanced yield (i.e. $[C/Fe] \geq +1.0$). Specifically, the yields were integrated over a Gaussian-like IMF centred on stars with a mass in the range $[10, 11, 12, 13.5, 15, 17, 20, 25, 30, 35, 40, 50, 75, 100 \, M_\odot]$, and with Gaussian width ranging from $0.025 - 0.5 \, M_\odot$ in steps of $0.025 \, M_\odot$. I adopted a piston explosion mechanism and the standard prescription of mixing for all models ($0.1 \, M_\odot$; see Section 4.7.2 for further details). Thus, there are a maximum of 20 models (due to the range of allowed Gaussian widths for the IMF) for each star of a given mass that explodes with a given final kinetic energy at infinity (which I henceforth refer to as the explosion energy$^1$). The distributions of $[C/Fe]$ values for these models are presented in Figure 6.1 where, instead of plotting all models for a given mass and explosion energy, I plot the median value from these models together with the associated dispersion. I have colour-coded the yields for each set of models by the explosion energy. It is readily apparent that the $[C/Fe]$ ratio depends on the progenitor mass for a given explosion energy — particularly for the lower mass progenitors.

A related observation from Figure 6.1 is that, for a given explosion energy, the fraction of

$^1$The total energy that is used to explode the star is equal to the star’s binding energy in addition to the final kinetic energy at infinity.
6.1 What fraction of DLAs are carbon-enhanced?

The C/Fe ratio is plotted for all Pop III models considered by Heger & Woosley (2010) that result in $[\text{C/Fe}] \geq +1.0$ (see text for further details). The symbols are colour-coded based on the final kinetic energy of the star’s yield, as per the legend provided. Note that if a Pop III star explodes with a final kinetic energy at infinity $E_\infty \gtrsim 3 \times 10^{51}$ erg, then according to these models, it will not produce a carbon-enhanced system.

Figure 6.1: The C/Fe ratio is plotted for all Pop III models considered by Heger & Woosley (2010) that result in $[\text{C/Fe}] \geq +1.0$ (see text for further details). The symbols are colour-coded based on the final kinetic energy of the star’s yield, as per the legend provided. Note that if a Pop III star explodes with a final kinetic energy at infinity $E_\infty \gtrsim 3 \times 10^{51}$ erg, then according to these models, it will not produce a carbon-enhanced system.

models that give rise to carbon-enhanced systems depends on the mass of the progenitor star. By integrating the fraction of carbon-enhanced models over a Salpeter-like IMF, one can obtain a rough estimate of the power law index from the observed fraction of CEMP DLAs, $f_{\text{CEMP}}$, provided that the “typical” explosion energy of metal-free stars is known. I therefore show in Figure 6.2 the dependence of $f_{\text{CEMP}}$ on the power law index for the primordial IMF, for a range of explosion energies.

However, even if we were to assume some typical explosion energy (such as the “standard” kinetic energy of $1.2 \times 10^{51}$ erg; see e.g. Woosley & Weaver 1995), $f_{\text{CEMP}}$ is a difficult quantity to measure in practice; Almost by definition, CEMP DLAs are those most likely to have strong saturated CII absorption lines. When the CII lines are saturated, it becomes very difficult to accurately measure the abundance of carbon (see the discussion in Section 2.3.1), and thus, carbon-enhanced systems may be easily overlooked. Indeed, very few DLAs are known where the carbon abundance can be accurately measured (see Table 4.11 for a full list).

Of course, since such DLAs will almost certainly later form a generation of Pop II stars (see Chapter 4), one could instead consider the fraction of CEMP stars in the halo of our Galaxy as a proxy for the fraction of CEMP DLAs. Again, this is hardly a straightforward task, since stars could exhibit carbon-enhancements from a variety of means (e.g. mass transfer from an asymptotic giant branch star), and may therefore not solely reflect the signature of Population III nucleosynthesis. Additional complications arise for the stellar sample, since we are only
Figure 6.2: The fraction of carbon-enhanced models is plotted as a function of the index for a Salpeter-like IMF ($\gamma_{\text{Sal}} = 1.35$ corresponds to a Salpeter IMF, represented by the vertical dashed line) for a range of explosion energies (colour-coded as in Figure 6.1). Black corresponds to a low explosion energy, whilst red corresponds to a high explosion energy. The grey band represents the current observational constraints on the fraction of carbon-enhanced systems.

probing the systems that formed stars in the halo of our Galaxy; the Galactic CEMP fraction may differ from the Universal CEMP fraction. Approximately, DLAs will likely provide a lower limit (due to the overlooked CEMP DLAs) and metal-poor stars will provide an upper limit (due to sample contamination). Thus, from the present sample of 10 DLAs there is one known CEMP DLA, which implies that $f_{\text{CEMP}} \geq 10\%$, whereas metal poor stars, on the other hand, suggest that $f_{\text{CEMP}} \leq 25\%$ (Lucatello et al., 2006). The range of possible values for the CEMP fraction is shown in Figure 6.2 as the grey band.

A possible way to circumvent the saturation problem for DLAs is to search for a diagnostic ratio that acts as a proxy for carbon-enhanced DLAs. For the reasons I will now discuss, one potential diagnostic that could be used to discriminate between CEMP and non-CEMP DLAs when the C II lines are saturated, is [N/Al]. In stars, N is synthesised in the H burning region via the CNO bi-cycle, and its production depends on the amount of C that is present. Thus, since C is enhanced, the production of N will be slightly enhanced if some C (which is synthesised in the He burning region), is brought into the H burning region, perhaps through rotation (see e.g. Hirschi 2007) or hot bottom burning (see e.g. Pagel 2009). The synthesis of Al depends on the neutron excess, which is low for metal-free stars (see e.g. Heger & Woosley 2002). In this picture, [N/Al] would thus be greatly enhanced in CEMP DLAs relative to their non-CEMP cousins, not only due to their enhanced N, but also due to the Al deficiency.

Such a diagnostic could prove to be quite valuable, since there exist a host of N I absorption lines of varying strength in DLAs. It is therefore likely that one could in principle find
6.1 What fraction of DLAs are carbon-enhanced?

Figure 6.3: The N/Al ratio is plotted for the CEMP DLA (green triangle) and ‘carbon-normal’ DLAs (red triangles) as a function of the O abundance. Upper limits on the N abundance are shown as the pink symbols with arrows. The dashed line represents the solar level of [N/Al].

Figure 6.4: The [N/Al] yield from the carbon-enhanced models of Heger & Woosley (2010), where the symbols have the same meaning as described above (cf. Figure 6.1). In virtually all cases, the [N/Al] ratio in these models is supersolar, with a lower plateau corresponding roughly to the value measured for the only known CEMP DLA.

To further illustrate the possibility of using the N/Al ratio to estimate the fraction of carbon-enhanced systems, I have plotted in Figure 6.4 the [N/Al] yields from the carbon-enhanced models of Heger & Woosley (2010), where the symbols have the same meaning as described above (cf. Figure 6.1). In virtually all cases, the [N/Al] ratio in these models is supersolar, with a lower plateau corresponding roughly to the value measured for the only known CEMP DLA. Again, the fraction of carbon-enhanced systems using this ‘unbiased’ technique is $f_{\text{CEMP}} = 10\%$. Once the statistics of metal-poor DLAs increases, it will be of interest to re-address this issue. Nevertheless, for the “standard” explosion energy of $1.2 \times 10^{51}$ erg (green line in Figure 6.2), the observed fraction of carbon-enhanced systems favours a power law IMF that is steeper than Salpeter. If, however, the “typical” explosion energy was larger by just 25% (i.e. $1.5 \times 10^{51}$ erg), the observed fraction of carbon-enhanced systems would imply an IMF that is shallower than Salpeter. It will be of interest to re-investigate this point again in the future,
Figure 6.4: The N/Al ratio is plotted for models of Population III nucleosynthesis where the [C/Fe] ratio is $\geq +1.0$ (as described earlier; cf. Figure 6.1). The dashed line represents the solar level of [N/Al]. Note that the models with [C/Fe] $\geq +1.0$ produce supersolar [N/Al].

Figure 6.3: The [N/Al] ratio is plotted for models of Population III nucleosynthesis where the [C/Fe] ratio is $\geq +1.0$ (as described earlier; cf. Figure 6.1). The dashed line represents the solar level of [N/Al]. Note that the models with [C/Fe] $\geq +1.0$ produce supersolar [N/Al].

when either the fraction of carbon-enhanced systems or the typical explosion energy of the first stars, is known to better accuracy.

### 6.2 Uncovering the nature of the first stars

The work presented in this thesis proposes that CEMP DLAs can be used to study the nucleosynthesis of the first stars. Whilst the non-CEMP DLAs might also contain the chemical signature from metal-free stars, it is less straightforward to decouple the Pop II enriched gas from the Pop III enriched gas; with CEMP DLAs, the delineation is simple.

Given the present sample of one CEMP DLA, it is difficult (perhaps impossible?) to constrain the parameters that characterise the first generation of stars, including the explosion energy, the degree of mixing, and ultimately, the form of the primordial IMF. Unfortunately, as demonstrated in Section 6.1, CEMP DLAs are biased to low energy explosions (cf. Figure 6.1). We therefore need to establish a simple means to remove this parameter from the analysis. One way to achieve this goal might be to consider the ratio of two Fe-peak elements. For example, previous work by Umeda & Nomoto (2002) suggests that high explosion energies for metal-free stars (i.e. hypernovae; $E_{\text{exp}} \sim 20 \times 10^{51}$ erg) can explain the large [Zn/Fe] ratio observed in the most metal-poor stars (see their Figure 14). Their [Zn/Fe] ratio, however, depends on the progenitor mass and hence on the IMF — which should be considered an unknown quantity for the moment.
Figure 6.5: Two diagnostic ratios are plotted for models of Population III nucleosynthesis where the $[\text{C}/\text{Fe}]$ ratio is $\geq +1.0$ (as described earlier; cf. Figure 6.1). The coloured lines indicate the different explosion energies (see text), where black line represents the lowest explosion energy and the red line represents the highest explosion energy. The dashed line represents the solar levels in all cases.

To circumvent this problem, in the left panel of Figure 6.5 I plot the abundance of $[\text{Ni}/\text{Fe}]$ (both Fe-peak elements) as a function of $[\text{Cr}/\text{Fe}]$ (also Fe-peak elements) for the sample of Pop III models discussed in Section 6.1 that deliver a carbon-enhanced signature relative to iron. For a given explosion energy, I have plotted a single line that represents the main ridge for all models that are considered here. Similarly, in the right panel, I plot $[\text{Ni}/\text{Fe}]$ versus $[\text{Zn}/\text{Fe}]$. For almost all models considered, one can uniquely determine the typical explosion energy for the stars that enriched a given system by considering the $[\text{Ni}/\text{Fe}]$ versus $[\text{Cr}/\text{Fe}]$ plot. For higher explosion energies, the $[\text{Cr}/\text{Fe}]$ ratio becomes degenerate, and the explosion energy is not unique when $[\text{Ni}/\text{Fe}] \lesssim -0.1$. In this case, one can instead use the right panel, where the higher explosion energies are well-separated. In this case, however, one will require spectra of very high $S/N$, since the Zn II lines are very weak and are considerably sub-solar relative to iron (which is already weak). The method of calculating the explosion energy from the $[\text{Ni}/\text{Fe}]$ versus $[\text{Cr}/\text{Fe}]$ diagnostic described above is feasible with the current generation of telescopes, provided that: (1) there is a bright background source ($m_r \sim 18$) that would enable one to efficiently obtain a high signal-to-noise ratio; (2) A DLA that is not too Fe-poor ($[\text{Fe}/\text{H}] \sim -2.5$); and (3) A DLA that has a relatively high H I column density (say, log $N$(H I)/cm$^{-2} \sim 20.7$).

For example, if $[\text{Ni}/\text{Fe}] = 0.0$ and $[\text{Cr}/\text{Fe}] = 0.0$, a system meeting the above criteria would require a signal-to-noise ratio per pixel of 40 and 50 to respectively detect the Ni and Cr absorption lines at the $3\sigma$ level (assuming the line profiles have a full width at half maximum of 5 km s$^{-1}$ corresponding to a Doppler parameter of $\sim 3$ km s$^{-1}$). Given the present telescope facilities and instruments (i.e. an 8m class telescope equipped with UVES), such a signal-to-noise ratio can be achieved with just $\sim 8$ one hour exposures for a spectral resolution of $\sim 8$ km s$^{-1}$ FWHM. Although not ideal, more favoured instrument setups are expected in the immediate
future; ESPRESSO\textsuperscript{2}, which is due to see first light on the VLT telescopes in 2016, will provide a spectral resolution of $\sim 2 \text{ km s}^{-1}$ FWHM, thus enabling one to fully resolve the intrinsically narrow absorption line profiles in the systems of interest. Furthermore, ESPRESSO can be fed with the light from all four VLT telescopes simultaneously, providing a collecting area that is equivalent to a 16 m diameter telescope. This setup will provide greater detail on the line profiles at an increased efficiency (in terms of total exposure time). The required exposure time will be further reduced when the next generation of extremely large telescopes come online, such as the European Extremely Large Telescope, which is scheduled to see first light shortly after 2020.

It is worth noting that in Figure 6.5, I have considered a model with relatively efficient mixing. In the event of less-efficient mixing, these diagnostic ratios will become further separated, therefore making it easier to distinguish between the different explosion energies. Nevertheless, before the explosion energy can be derived with certainty, one needs to establish the degree of mixing involved. Ideally, we require an element ratio that is different for varying degrees of mixing, but for all cases is invariant with the explosion energy and is independent of progenitor mass (and hence the IMF). In practice, given the small handful of elements available from a typical metal-poor DLA, such an ideal diagnostic is difficult to come by. Thus, whilst we may not be able to infer the degree of mixing for a given case, we can instead statistically determine the amount of mixing that applies to the general population of metal-free stars by considering a ratio that is independent of the IMF. For example, in Figure 6.4, one can see that the $[\text{N}/\text{Al}]$ ratio is essentially independent of progenitor mass, although clearly dependent on the explosion energy.

The distribution of these models is presented in the top panel of Figure 6.6 as the thick solid black line. Also shown, are the distributions for the individual explosion energies (with the same colour scheme as in Figure 6.4). In the bottom panel of Figure 6.6, I instead plot the distributions for the case of less-efficient mixing. In either mixing scenario, roughly the same amount of $^{14}\text{N}$ is ejected (see Figure 7 from Heger & Woosley 2010), since N is primarily synthesised in the outer regions of the star. Thus, the final $[\text{N}/\text{Al}]$ ratio is largely driven by the amount of Al ejected. Consider first the models with low explosion energy. Less mixing in these models will result in less Al being mixed to the outer regions where it can be ejected. As a consequence, $[\text{N}/\text{Al}]$ shifts to higher values for less mixing in low energy explosions. Higher energy explosions, on the other hand, can remove Al without the need for mixing, due to their higher energy. However, less mixing in models with a high explosion energy causes less Fe to be mixed outwards and thus ejected. This results in a larger percentage of CEMP models with high explosion energies. Since models with a high explosion energy are able to eject more of the Al, the distributions of $[\text{N}/\text{Al}]$ move to lower values for high energy explosions. Thus, the $[\text{N}/\text{Al}]$ distribution becomes less concentrated when there is less mixing.

\textsuperscript{2}An acronym for Echelle SPectrograph for Rocky Exoplanet and Stable Spectroscopic Observations
In fact, the [N/Al] ratio alone may provide information for both the explosion energy and the mixing that gives rise to the carbon-enhanced systems; the width of the distribution provides the mixing, and the centre of the distribution (once the mixing is known) estimates the explosion energy. From here, one could then determine the characteristic mass scale from a plot such as that presented in Figure 6.1 or 6.2, where the explosion energy and mixing provide an (approximate) estimate for the characteristic mass scale. Although, to achieve the accuracy required to measure the width and centroid of this distribution, one would need to determine the [N/Al] ratio with an uncertainty \( \lesssim 0.1 \) dex, which is certainly plausible given the currently available instrumentation and techniques. The largest uncertainty, however, is likely the result of uncertain ionisation corrections to the [N/Al] ratio. As I have shown previously in Figures 4.9.

**Figure 6.6:** The distribution of [N/Al] is plotted for models of Population III nucleosynthesis where the [C/Fe] ratio is \( \geq +1.0 \) (as described earlier; cf. Figure 6.1) for the standard (efficient) mixing case (thick black histogram; top panel). Also shown are the distributions of [N/Al] for each explosion energy, represented by the same colour-code as used in Figure 6.1. The vertical dashed lines indicate the median values for each distribution. I have also plotted the same distributions for the case of less-efficient mixing (lower panel). In the case of less mixing, the distribution of [N/Al] ratios becomes broader, and the median values for each explosion energy become more separated.
and 5.3, the differential correction for N/Al can be as high as $\sim 0.4$ dex at a number density of $10^{-3}$ cm$^{-3}$. Such low densities are, however, unlikely to occur in these DLAs; more typical densities of $n \sim 0.1 - 1 \text{ cm}^{-3}$ are expected, which correspond to ionisation corrections less than $\lesssim 0.1$ dex. Indeed, by not applying these ionisation corrections, there might exist some residual scatter, or perhaps a shift in the centroid of this distribution. This bias should, however, be negligible compared with the $\sim 1.0$ dex scatter seen for the different degrees of mixing.

At this point, it is worth bearing in mind that by studying CEMP systems, we are likely not biased to different progenitor masses (i.e. any progenitor mass can produce a carbon-enhanced system, see Figure 6.1); the carbon-enhanced models considered here are not carbon-enhanced due to the progenitor mass, they are carbon-enhanced due to the properties of the explosion. It is therefore expected that by using these objects, the primordial IMF is still being sampled in an unbiased fashion. In other words, the characteristic mass scale for the general population of Pop III stars (i.e. independent of their C/Fe yield) is likely independent of the mixing and possibly the explosion energy (or more specifically, the final kinetic energy of the ejected mass). Of course, these conclusions depend on the adopted models and are thus based on the still uncertain explosion mechanism that operates in CCSNe. Once there is a better handle on this problem, it will be of interest to re-examine the diagnostic ratios that can provide hints into the mass function of the first stars.

The above considerations outline a simple and straightforward method to study the nature of the first generation of stars. I have shown that if one can find the first metals before they are incorporated into the second generation of stars, then measuring the chemical composition is less complicated and more reliable than alternative means. A selection of the most-metal poor DLAs is proposed to be an easily identifiable missing link between the first and second generation stars, and thus harbour the chemical signature from metal-free stars. By building a small sample of such systems, one might hope to unveil the nature and evolution of the elusive first generation of stars.
In this supplementary Chapter, I employ the Union compilation of Type Ia supernovae (SNe) with a maximum likelihood analysis to search for a dark energy dipole. To approach this problem, I present a simple, computationally efficient, and largely model independent method. This technique involves weighting each SN by its observed error estimate, so that poorly measured SNe that deviate substantially from the Hubble law do not produce spurious results. I find, with very low significance, a dipole in the cosmic acceleration directed roughly towards the cosmic microwave background (CMB) dipole, but this is almost certainly coincidental.
7.1 Introduction

The foundations for the Friedmann-Lemaître-Robertson-Walker (FLRW) cosmological model rest on the assumptions of isotropy and homogeneity, both of which seem plausible on average. However, given the discovery of cosmic acceleration (Riess et al., 1998; Perlmutter et al., 1999), the FLRW model must attribute this observation to a cosmological constant (dark energy) in Einstein’s equations, a constant that still lacks an acceptable physical interpretation (see Frieman, Turner & Huterer 2008 for a review).

To alleviate these concerns, numerous authors have investigated possible inhomogeneous, and void models to account for the observed acceleration (see Enqvist 2008 for a review, but also Alnes, Amarzguioui & Grøn 2006 and references therein). The typical approach is to employ the exact Lemaître-Tolman-Bondi solutions for a radially inhomogeneous, spherically symmetric universe, and require that we must be almost centrally located in a large-scale void. However, a need to violate the Copernican principle has not yet been evidenced. Moreover, many of these models seem unlikely (see Ishibashi & Wald 2006; Moss, Zibin, & Scott 2010).

A more recent development by Wiltshire (2007a), suggests that cosmic acceleration can be explained in general relativity, by considering the differences in the quasi-local gravitational energy between observers in bound systems (clusters of galaxies in bubble walls and filaments) and those in a freely expanding space (volume-averaged observers), resulting in a cumulative time dilation effect due to the differences in their clock rates (see also Wiltshire 2007b). Using the Gold SNe sample (Riess et al., 2007), Leith, Ng & Wiltshire (2008) demonstrate that this model provides a possible alternative to dark energy, although this model fails to provide a consistent fit to the Union (Kowalski et al., 2008) and Constitution (Hicken et al., 2009) SNe samples (Kwan, Francis & Lewis, 2009), contrary to the current concordant cosmology ($\Lambda$CDM), whereby all three SNe samples are fit by a consistent set of cosmological parameters.

Other authors have investigated the possibility that dark energy is anisotropic (see Cooray, Holz, & Caldwell 2010 and references therein). Cooray, Holz, & Caldwell (2010) present a viable method to model dark energy inhomogeneities as a power spectrum of luminosity fluctuations. However, with such optimistic objectives, these studies will require a dedicated survey to search for SNe, such as the Dark Energy Survey (Frieman et al., 2005).

One can, however, apply a simple method to measure the anisotropy of dark energy or, alternatively, search for a systematic directional dependence in the SNe data (Kolatt & Lahav, 2001; Schwarz & Weinhorst, 2007; Gupta, Saini, & Laskar, 2008). This involves dividing the data into two hemispheres and searching for the strongest directional dependence, which these authors find to have a significance of about 1-2$\sigma$ for previous compilations of SNe.

In this Chapter I investigate the employability of the current SNe dataset to test for the presence of a dark energy dipole. In Section 7.2 I detail the SN sample used in my analysis, and then formulate the dipolar model in Section 7.3. In Section 7.4 I present the results of the dipolar modelling, and summarise my conclusions in Section 7.5. Throughout I adopt the
7.2 The Type Ia Supernova Sample

For the present analysis, I employ the Union sample of SNe, compiled by Kowalski et al. (2008). The Union compilation contains 307 identically processed SNe, drawn from a heterogeneous sample of 414. For the present work, it is noteworthy that Kowalski et al. (2008) corrected their data for the CMB dipole, and rejected SNe with CMB-centric redshifts $z < 0.015$. The Galactic coordinates for all SNe in the Union sample were obtained from the NASA/IPAC Extragalactic Database (NED). Indeed, as illustrated in Figure 7.1, the Union SNe compilation are drawn from a spatially uniform sample, except of course the trivial lack of SNe observed in the Galactic plane.

It should be noted that SN compilations are not exempt from selection effects, such as Malmquist bias, and should be modelled through Monte Carlo simulations (see for example Kessler et al. 2009). These considerations are beyond the scope of the current work, and in any case, should not have an angular dependence.

The Union SN compilation was recently updated by Hicken et al. (2009) to include additional low redshift SNe. I have decided not to include these additional data for the following two reasons: First, the signature of an anisotropic acceleration of the Universe is only detectable at higher redshifts, where the effect of a cosmological constant is more exposed, and the motion cosmological parameters derived by Dunkley et al. (2009); $H_0 = 71.9$ km s$^{-1}$ Mpc$^{-1}$, $\Omega_m = 0.258$, and $\Omega_\Lambda = 0.742$.

7.2 The Type Ia Supernova Sample

For the present analysis, I employ the Union sample of SNe, compiled by Kowalski et al. (2008). The Union compilation contains 307 identically processed SNe, drawn from a heterogeneous sample of 414. For the present work, it is noteworthy that Kowalski et al. (2008) corrected their data for the CMB dipole, and rejected SNe with CMB-centric redshifts $z < 0.015$. The Galactic coordinates for all SNe in the Union sample were obtained from the NASA/IPAC Extragalactic Database (NED). Indeed, as illustrated in Figure 7.1, the Union SNe compilation are drawn from a spatially uniform sample, except of course the trivial lack of SNe observed in the Galactic plane.

It should be noted that SN compilations are not exempt from selection effects, such as Malmquist bias, and should be modelled through Monte Carlo simulations (see for example Kessler et al. 2009). These considerations are beyond the scope of the current work, and in any case, should not have an angular dependence.

The Union SN compilation was recently updated by Hicken et al. (2009) to include additional low redshift SNe. I have decided not to include these additional data for the following two reasons: First, the signature of an anisotropic acceleration of the Universe is only detectable at higher redshifts, where the effect of a cosmological constant is more exposed, and the motion

---

1 Available from: http://supernova.lbl.gov/Union/
2 Kowalski et al. (2008) find that their results do not depend significantly on the adopted redshift cutoff.
3 http://nedwww.ipac.caltech.edu/
of the local group is less of a contaminant. Therefore a low redshift sample should not affect the final result. Moreover, for this reason, I opted to include only the Union SNe with \( z > 0.2 \), resulting in a total sample of 250. Second, the more recent data are processed using different routines, which could introduce a bias if I was to blindly combine both samples. With the goal of the present work borne in mind, I restrict my analysis to the Union compilation.

### 7.3 Analysis

As outlined by (Kolatt & Lahav, 2001), a number of sources could produce deviations from isotropy in the SN data. From the results of their study, these authors also raised concerns about the relatively small number of SNe, spread inhomogeneously across the sky, with differing and uncertain errors. I therefore seek a method that is not biased by the clustering of the SN data, whilst taking into consideration the individual uncertainties of the SNe. Bearing in mind these characteristics, I employ a maximum likelihood strategy (unbiased to clustering) and devise a method that weights the contribution of each SN based on its corresponding error measurement.

#### 7.3.1 The Hubble Deviation

Before proceeding with a quantitative analysis of the Union sample, it is instructive to inspect whether any obvious preferred deviations from the Hubble law exist in the employed dataset. Following Jha, Riess & Kirshner (2007), the deviation from the Hubble law is given by

\[
\frac{\delta H}{H} = \frac{H_0 d_L - H_0 d_{SN}}{H_0 d_{SN}}
\]

(7.1)

where \( d_L \) is the luminosity distance in a \( k = 0 \) universe,

\[
d_L = \left( \frac{1+z}{H_0} \right) \int_0^z \left[ \Omega_\Lambda + \Omega_M (1+z')^3 \right]^{-1/2} \, dz'
\]

(7.2)

and \( d_{SN} \) can be calculated from the observed distance modulus, \( \mu = 5 \log d_{SN} + 25 \). Figure 7.2(a) illustrates that the data are very slightly skewed; more SNe are above the Hubble law (i.e. with negative Hubble deviation) than below (positive Hubble deviation). Moreover, perhaps unsurprisingly, there are also a greater number of SNe that lie above the Hubble law when considering the subsample of SNe that exhibit the largest Hubble deviations (\(| \delta H / H | > 0.15 \)), represented in Figure 7.1 and 7.2 by the crosses and open circles. The entire SNe sample are presented on a Hubble plot in Figure 7.2(b). The fact that all of the largest deviations are exhibited by SNe at \( z > 0.2 \), justifies my decision not to include SNe with \( z < 0.2 \).

The spatial distribution of the entire Union sample of SNe is presented in Figure 7.1, where the symbols correspond to those in Figure 7.2. Indeed, from this simple demonstration, an obvious dipole is not apparent.
7.3.2 Maximum Likelihood Strategy

I consider a $k = 0$ universe, and compare two models with and without an acceleration term. In essence, the result should not be affected by the model chosen to include the acceleration term, provided that the model represents the centroid of the data at a given redshift (i.e. selecting the $\Lambda$CDM model is sufficient). I define $\delta(\Lambda) = d(\Lambda) - d(\Lambda = 0)$ to be the luminosity distance deviation of an isotropically accelerating universe, $d(\Lambda)$, from a non-accelerating universe, $d(\Lambda = 0)$. The luminosity distance one would measure in a universe with an anisotropic acceleration $d(\Lambda_A)$, can then be simply described by

$$d(\Lambda_A) = d(\Lambda) + \delta(\Lambda) X$$

(7.3)

where $X$ varies over a range, say $\pm L$, and provides a measure for the degree of anisotropic acceleration. For example, if $X = 0$, one simply recovers $d(\Lambda)$, and if $X = -1$, one is left with $d(\Lambda = 0)$. Introducing a unit vector $\hat{l}_i$ that points in the direction of the $i^{th}$ SN, and a vector $\vec{L}$, that points in the direction of ‘preferred’ acceleration, I can suitably choose $X = \hat{l}_i \cdot \vec{L}$, thus
Eq. 7.3 becomes
\[ \delta_{M,i} = D(z_i)(1 + \hat{\ell}_i \cdot \vec{L}) \]  
(7.4)

with,
\[ \delta_{M,i} = \frac{d(\Lambda) - d(\Lambda = 0)}{d(\Lambda = 0)} \]  
(7.5)

\[ D(z_i) = \frac{d(\Lambda) - d(\Lambda = 0)}{d(\Lambda = 0)} \]  
(7.6)

where \( D(z_i) \) is the ‘deviation parameter’ for the \( i \)th SN with redshift \( z_i \). Using this formalism, I introduce the measured quantity \( \delta_i \) (recall that \( d_{SN} \) is the measured luminosity distance to a given SN),
\[ \delta_i = \frac{d_{SN} - d(\Lambda = 0)}{d(\Lambda = 0)} \]  
(7.7)

The Likelihood Function

The dipolar model (Eq. 7.4) forms the basis for the present analysis, and is employed using a maximum likelihood strategy. As the observations are not uniformly distributed around the sky, I pose the question, given the directions \( \hat{\ell}_i \) of the observed SNe, what is the most probable vector \( \vec{L} \) that gives an anisotropy of the form in Eq. 7.4? Non-uniform sky coverage will therefore only result in a larger error ellipsoid for \( \vec{L} \).

Within this framework, poorly measured SNe that deviate substantially from the concordance model will produce spurious results. I therefore weight each SN according to its observational error. Suppose the measurements \( \delta_i \) come from true values \( \Delta_i \) scattered by observational errors \( \sigma_i \). Then, given an observation \( \delta_i \), the probability of it arising from a true value \( \Delta_i \) is
\[ Pr(\delta_i | \Delta_i) = \frac{1}{\sqrt{2\pi \sigma_i}} \exp \left( -\frac{(\delta_i - \Delta_i)^2}{2\sigma_i^2} \right) \]  
(7.8)

where \( \sigma_i \) is the observers error estimate for the \( i \)th SN. Similarly, the probability that the true \( \Delta_i \) arise from a scatter \( \sigma \) about the proposed dipolar model (Eq. 7.4), is
\[ Pr(\Delta_i | \delta_{M,i}) = \frac{1}{\sqrt{2\pi \sigma}} \exp \left( -\frac{(\Delta_i - \delta_{M,i})^2}{2\sigma^2} \right) \]  
(7.9)

The probability, therefore, of measuring the \( \delta_i \) given the model is found by integrating over all
true values

\[ Pr(\delta \mid \delta_{M,i}) = \int_{-\infty}^{\infty} Pr(\delta \mid \Delta_i) \cdot Pr(\Delta_i \mid \delta_{M,i}) \ d\Delta_i \]

\[ Pr(\delta \mid \delta_{M,i}) = \frac{1}{\sqrt{2\pi(\sigma^2 + \sigma_i^2)}} \exp \left( -\frac{(\delta_i - \delta_{M,i})^2}{2(\sigma^2 + \sigma_i^2)} \right) \] (7.10)

The log-likelihood function is then given by

\[ \mathcal{L} = \ln \left[ \prod_i Pr(\delta_i \mid \delta_{M,i}) \right] \] (7.11)

**Maximum Likelihood**

Substitution of the dipolar model into Eq. 7.11, and maximising the likelihood over all choices of the model parameters \( \vec{L} \) and \( \sigma \), \( \partial \mathcal{L} / \partial \vec{L} = 0 \) yields an equation of the form

\[ \mathbf{A} \cdot \vec{L} = \vec{V} \] (7.12)

where \( \mathbf{A} \) (a rank 2 tensor), and \( \vec{V} \) are given by

\[ \mathbf{A} = \sum_i \frac{D(z_i)^2}{\sigma^2 + \sigma_i^2} \hat{l}_i \hat{l}_i \] (7.13)

\[ \vec{V} = \sum_i \frac{(\delta_i - D(z_i)) D(z_i)}{\sigma^2 + \sigma_i^2} \hat{l}_i \] (7.14)

\[ \partial \mathcal{L} / \partial \sigma = 0 \] gives

\[ \sum_i \frac{1}{\sigma^2 + \sigma_i^2} = \sum_i \frac{(\delta_i - \delta_{M,i})^2}{(\sigma^2 + \sigma_i^2)^2} \] (7.15)

One must now simultaneously solve the coupled non-linear Eqs. 7.12 and 7.15. The most efficient pathway to a solution is to construct a set of \( \sigma \) values. Using Eq. 7.12, one can invert \( \mathbf{A} \), and find \( \vec{L} \) for each \( \sigma \). One can then use Eq. 7.15 to solve for the \( \sigma \) value that maximises \( \mathcal{L} \).

**Other Solution Pathways**

The non-linearity of the above equations is introduced through the weighting of each SN by its associated error. If one weights all SNe equally (i.e. \( \sigma_i = \text{constant} \) for all \( i \), Eqs. 7.13-7.15 then take the form

\[ \mathbf{A} = \sum_i D(z_i)^2 \hat{l}_i \hat{l}_i \] (7.16)
\[ \vec{V} = \sum_i (\delta_i - D(z_i)) D(z_i) \hat{l}_i \]  
\[ \sigma^2 = \frac{1}{N} \sum_i (\delta_i - \delta_{M,i})^2 \]  
(7.17)

where \( N \) is the total number of SNe in the sample. One can now invert \( A \) and uniquely solve for \( \vec{L} \), and then use Eq. 7.18 to solve for the \( \sigma \) that maximises the likelihood, \( \mathcal{L} \).

### 7.4 Results and Discussion

In this section I present the results from the maximum likelihood analysis described above, and introduce a method to quantify the significance of the derived parameter values. Before proceeding, it is noteworthy that the adopted maximum likelihood strategy detailed in Section 7.3.2 requires all SNe coordinates to be converted from Galactic to Cartesian coordinates\(^4\).

#### 7.4.1 Significance of the Dipolar Models

To calculate the significance of the dipolar model, one can ask, what is the probability the vector \( \vec{L} \), that maximises the likelihood, is true? The covariance matrix for \( \vec{L} \) is given by the inverse of Eq. 7.13 (i.e. \( A^{-1} \)), or, for the solution pathway outlined in Section 7.3.2, \( \sigma^2 A^{-1} \). Given the eigenvalues \( (\lambda_1, \lambda_2, \lambda_3) \) and corresponding eigenvectors \( (\hat{e}_1, \hat{e}_2, \hat{e}_3) \) of the covariance matrix, the three semi-principal axes of the error ellipsoid for \( \vec{L} \) are \( \hat{s}_{1,2,3} = \sqrt{\lambda_{1,2,3}} \hat{e}_{1,2,3} \). The density of probability associated with the vector \( \vec{L} \) is therefore of the form

\[ p(\vec{L}) = \frac{1}{\sqrt{(2\pi)^3 \lambda_1 \lambda_2 \lambda_3}} \exp \left( -\frac{e_1^2}{2\lambda_1} - \frac{e_2^2}{2\lambda_2} - \frac{e_3^2}{2\lambda_3} \right) \]  
(7.19)

where \( e_1, e_2 \) and \( e_3 \) are the Cartesian coordinate axes in the direction of the principal axes of the error ellipsoid. The length of the vector \( \vec{L} \) enclosed by the error ellipsoid is given by

\[ L_{\text{err}} = [(\vec{L} \cdot \hat{s}_1)^2 + (\vec{L} \cdot \hat{s}_2)^2 + (\vec{L} \cdot \hat{s}_3)^2]^{1/2} \]  
(7.20)

I now introduce the quantity \( \mu = L_{\text{err}}/|\vec{L}| \), which defines the fraction of \( \vec{L} \) enclosed by the error ellipsoid. Suppose one now expands (or contracts) the error ellipsoid such that it contains a volume \( V_E \), that just encompasses \( \vec{L} \). The probability that the derived \( \vec{L} \) is true, is therefore given by

\[ Pr(\vec{L}) = \int_{V_E} p(\vec{L}) \, dV_E \]  
(7.21)

\(^4\)The \( x \) and \( z \) Cartesian axes are respectively directed towards the Galactic centre and North pole
Table 7.1: Parameter fitting results

| Model\textsuperscript{a} | $|\vec{L}| \times 10^2$ | $l^\dagger$ | $b^\dagger$ | $\sigma \times 10^2$ | $\vec{L}^\dagger$ | $L_{err}/|\vec{L}|$ | $Pr(\vec{L})$ |
|---------------------------|----------------------|------------|------------|------------------|----------------|----------------|-------------|
| Sample 1                  | 7.14                 | 287°       | 46°        | 0.00             | 105.5          | 1.04           | 18% (0.23$\sigma$) |
| Sample 2                  | 3.90                 | 258°       | 9°         | 0.00             | 82.9           | 3.44           | 0.6% (0.008$\sigma$) |
| Sample 3                  | 13.8                 | 309°       | 43°        | 3.52             | 23.3           | 0.83           | 31% (0.39$\sigma$) |
| Sample 4                  | 10.0                 | 118°       | 25°        | 14.2             | 97.1           | 1.03           | 18% (0.23$\sigma$) |

\textsuperscript{a}Sample 1, 2, 3: Weighted Dipolar Model with respectively $z > 0.2$, $0.2 < z < 0.56$, and $z > 0.56$. Sample 4: Higher Quality List. $\dagger$Galactic Coordinates. $\ddagger$Given by Eq. 7.11.

which after converting to spherical coordinates yields

$$Pr(\vec{L}) = \frac{4\pi}{(\pi)^{3/2}} \int_0^{1/\mu \sqrt{2}} u^2 \exp(-u^2) \, du$$

$$Pr(\vec{L}) = \text{erf}(1/\mu \sqrt{2}) - \sqrt{\frac{2}{\mu^2 \pi}} \exp(-1/2\mu^2)$$  \hspace{1cm} (7.22)

7.4.2 The Weighted Dipolar Model

I now consider the weighted dipolar model; weighting all SNe according to their measured errors, and implement the solution strategy outlined in Section 7.3.2. The model parameters that maximise the likelihood function are presented in Table 7.1 (Sample 1), where the components of $\vec{L}$ are now converted back into Galactic coordinates. For this case, $\sigma^2$ is negligible compared to the minimum $\sigma_i^2$ in the sample (refer to Eq. 7.10). The fitting results are also illustrated in Figure 7.3(a), where the data are averaged in 10 equally spaced bins, and the error bars correspond to the standard error in the mean of each bin. Bins containing $\leq 2$ points are not shown.

A striking observation of the results from the weighted dipolar model, is the similarity in the direction of $\vec{L}$ to the CMB dipole: $(l, b) = (263.99^\circ \pm 0.14^\circ, 48.26^\circ \pm 0.03^\circ)$ (Hinshaw et al., 2009). Although there is a relatively minor significance, this coincidence is certainly worth further scrutiny.

To investigate whether the higher or lower redshift SNe, or both, favour this direction, I divide the SNe into two subsamples, each containing 125 SNe, and then run the weighted dipolar model routine on the two subsamples. This division corresponds to a redshift $z = 0.56$. The results are again presented in Table 7.1, and illustrated in Figure 7.3(b) and 7.3(c) for the two subsamples $0.2 < z < 0.56$ (Sample 2) and $z > 0.56$ (Sample 3) respectively. As can be appreciated from Figure 7.3(b), and the final column of Table 7.1, the lower redshift subsample does not resemble the anisotropic model at all. The higher redshift subsample presented in Figure 7.3(c), on the other hand, seems more consistent with the results of Sample 1.

In order to test whether the Union compilation exhibit a dipolar acceleration, I plot the weighted data as a function of $\vec{l}_i \cdot \vec{L}$ (i.e. the cosine of the angle between the $i$th SN and the
Figure 7.3: For all figures the y-axis corresponds to the deviation from an isotropic model, and the x-axis corresponds to the cosine of the angle between the $i^{th}$ SN and the 'preferred' direction $\vec{L}$, weighted by the deviation parameter and the individual errors, so that poorly measured SNe contribute less, as required by the dipolar model. Thus, the solid horizontal line corresponds to an isotropic model, and the dotted line has a gradient of $|\vec{L}|$. For a perfect fit to the dipolar model, the data should lie along this dotted line. The data are averaged in 10 equally spaced bins, where the errors represent the standard error in the mean of the given bin. The domain is set to be the maximum abscissa of the unbinned data. (a) Weighted dipolar model (Sample 1) $z > 0.2$. (b) Weighted dipolar model (Sample 2) $0.2 < z < 0.56$. (c) Weighted dipolar model (Sample 3) $z > 0.56$. (d) Higher Quality List (Sample 4).

‘preferred direction’) in Figure 7.4(a)-(c). Should a preferred direction exist, one would expect the data to show a general trend from the bottom-left quadrant to the top-right quadrant, a feature that is not observed in these samples. One could tentatively claim that this is true for Sample 3 (in Figure 7.4(c)), but the significance is very low.

7.4.3 The Higher Quality List

In Section 7.3.2, an alternative linear method was described to solve for the model parameters, however, as mentioned previously, one must be cautious when assigning each SN measurement an equal weighting, as poorly measured SNe that deviate substantially from the Hubble law could produce spurious results. These concerns should be alleviated by compiling a ‘Higher Quality List’, containing only the best quality SNe (Sample 4). I therefore restrict the sample to contain only the best measured SNe; SNe with uncertainties in magnitude $> 0.35$ are rejected. The distribution of accepted and rejected SNe are presented in Figure 7.5(a). The Hubble diagram for the Higher Quality List is shown in Figure 7.5(b).

One would expect the results of the weighted dipolar model (Sample 1) to be consistent with the Higher Quality List, however, this does not seem to be the case. The direction of $\vec{L}$ has
7.5 Conclusions

The acceleration of the Universe appears to be equal in all directions. The anisotropy I find is a 14% ± 12% increase toward \((l, b) = (309°, 43°)\) and a corresponding decrease in the opposite direction \((l, b) = (129°, -43°)\), but the effect is only apparent in the higher redshift group with \(z > 0.56\). These are those that exhibit the largest net acceleration. The direction of this weak anisotropy is only 31° from that of the microwave dipole as seen from the Sun, \((l, b) = (264°, 48.3°)\), or 17° from that of the dipole observed in the CMB frame of the elliptical galaxies sample with \(v < 2000\, \text{km s}^{-1}\), \((l, b) = (311°, 26°)\) (Lynden-Bell et al., 1988). The Union SN redshifts are corrected to the CMB system of rest, so their zero point cannot cause the effect.

It is noteworthy that both Kolatt & Lahav (2001) and Gupta, Saini, & Laskar (2008) find a directional-dependent systematic in the SNe data of a significance similar to that found here,
although the directions are somewhat different. The present dataset cannot exclude a 26\% (i.e. 14+12) increase in the acceleration toward $(l, b) = (309^\circ, 43^\circ)$ and together with a 26\% decrease in the opposite direction. More high accuracy $z > 0.5$ data from directions within 45\° of $(l, b) = (309^\circ, 43^\circ)$ or $(l, b) = (129^\circ, -43^\circ)$ are needed to eliminate any possible anisotropy.

### 7.6 Future Work

The analytic maximum likelihood method employed in this Chapter has the advantage of being invariant to the clustering of data points, whilst taking into account the uncertainty in the distance measures. However, rather than calculating a dipole in the distance measures, as done here, one could similarly apply this method to calculate a dipole in velocity (or equivalently, in redshift), which is perhaps more intuitive than a dipole in luminosity distance. Whilst this will require a non-analytic approach, one can then simply test how the magnitude of this dipole varies as a function of redshift. This can be achieved by numerically converting the luminosity distance.
distance for each type Ia supernova into an “expected redshift” using the luminosity distance relation (Eq. 7.2). By then computing the deviations between this expected redshift and the measured redshift, the velocity dipole can be calculated. The significance of the dipole is then roughly found, for a given redshift bin, by dividing the dispersion in the deviations by the typical uncertainty of the data. At this point it is noteworthy that such an approach only tests for a systematic directional dependence in the supernovae data; it does not attribute the dipole to a physical effect.

The above method will require a larger sample of supernovae than that analysed in this Chapter in order to investigate the redshift evolution of a dipole. Recently, the Union2 compilation was released (Amanullah et al., 2010), which now contains 557 type Ia supernovae. Aside from investigating the presence of a dipole, these new data could provide interesting contraints on the potential evolution of a dipole with redshift.

Finally, the fact that a dipole is possibly observed in the higher redshift (as opposed to the lower redshift) supernovae analysed here, begs the question: At what point does the Universe expand with the Hubble flow? Suppose one uses the supernovae data without correcting for the CMB dipole, and then measures the velocity dipole for all supernovae in a given redshift shell. At some redshift, the data will exhibit no dipole, thus indicating that these supernovae are moving with the Hubble flow. If one was to now calibrate the zero-point of the type Ia supernovae based on the higher redshift sample (where all SNe are expected to be moving with the Hubble flow), one could test how similar the zero-point of the CMB dipole is to that of the Hubble flow. Future large compilations of SNe, such as those from the Dark Energy Survey (Frieman et al., 2005, scheduled to begin at the end of 2011), will provide a supernovae sample that is large enough to answer such questions.
APPENDICES
The adopted Solar abundance scale

Throughout this thesis I have used the recently updated solar abundance scale proposed by Asplund et al. (2009) as a baseline to compare my element abundances. In Table A.1, I provide both the photospheric and meteoritic abundances from this solar scale, including the solar abundances that I have adopted in the penultimate column of Table A.1. The final column of this table indicates that I have adopted either the photospheric abundance (P), the meteoritic abundance (M), or the average of the two (A), as suggested by Lodders, Plame, & Gail (2009).

<table>
<thead>
<tr>
<th>X</th>
<th>Photo / Meteor</th>
<th>Adopted</th>
<th>X</th>
<th>Photo / Meteor</th>
<th>Adopted</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>log(X/H)⊙</td>
<td>log(X/H)⊙</td>
<td></td>
<td>log(X/H)⊙</td>
<td>log(X/H)⊙</td>
</tr>
<tr>
<td>H</td>
<td>0.00</td>
<td>0.00</td>
<td>S</td>
<td>−4.88</td>
<td>−4.85</td>
</tr>
<tr>
<td>He</td>
<td>−1.07</td>
<td>−10.71</td>
<td>P</td>
<td>−6.50</td>
<td>−6.77</td>
</tr>
<tr>
<td>Li</td>
<td>−10.95</td>
<td>−8.74</td>
<td>Cl</td>
<td>−6.50</td>
<td>−12.50</td>
</tr>
<tr>
<td>Be</td>
<td>−10.62</td>
<td>−10.70</td>
<td>Ar</td>
<td>−5.60</td>
<td>−12.50</td>
</tr>
<tr>
<td>B</td>
<td>−9.30</td>
<td>−9.21</td>
<td>K</td>
<td>−6.97</td>
<td>−6.92</td>
</tr>
<tr>
<td>C</td>
<td>−3.57</td>
<td>−4.61</td>
<td>Ca</td>
<td>−5.66</td>
<td>−5.71</td>
</tr>
<tr>
<td>N</td>
<td>−4.17</td>
<td>−5.74</td>
<td>Sc</td>
<td>−8.85</td>
<td>−8.95</td>
</tr>
<tr>
<td>O</td>
<td>−3.31</td>
<td>−3.60</td>
<td>Ti</td>
<td>−7.05</td>
<td>−7.09</td>
</tr>
<tr>
<td>F</td>
<td>−7.44</td>
<td>−7.58</td>
<td>V</td>
<td>−8.07</td>
<td>−8.04</td>
</tr>
<tr>
<td>Ne</td>
<td>−4.07</td>
<td>−13.12</td>
<td>Cr</td>
<td>−6.36</td>
<td>−6.36</td>
</tr>
<tr>
<td>Na</td>
<td>−5.76</td>
<td>−5.73</td>
<td>Mn</td>
<td>−6.57</td>
<td>−6.52</td>
</tr>
<tr>
<td>Mg</td>
<td>−4.40</td>
<td>−4.47</td>
<td>Fe</td>
<td>−4.50</td>
<td>−4.55</td>
</tr>
<tr>
<td>Al</td>
<td>−5.55</td>
<td>−5.57</td>
<td>Co</td>
<td>−7.01</td>
<td>−7.13</td>
</tr>
<tr>
<td>Si</td>
<td>−4.49</td>
<td>−4.49</td>
<td>Ni</td>
<td>−5.78</td>
<td>−5.80</td>
</tr>
<tr>
<td>P</td>
<td>−6.59</td>
<td>−6.57</td>
<td>Cu</td>
<td>−7.81</td>
<td>−7.75</td>
</tr>
</tbody>
</table>

Table A.1: Adopted Solar Abundances
Column densities for VMP DLAs

To facilitate comparison with future data sets, Table B.1 lists the ion column densities for the VMP DLAs analysed in Chapter 4. The corresponding element abundances, based on the solar abundance scale listed in Appendix A, are collected in Table 4.11.
<table>
<thead>
<tr>
<th>QSO</th>
<th>$z_{\text{abs}}$</th>
<th>$\log N(\text{H} I)$ (cm$^{-2}$)</th>
<th>$\log N(\text{C} II)$ (cm$^{-2}$)</th>
<th>$\log N(\text{N} I)$ (cm$^{-2}$)</th>
<th>$\log N(\text{O} I)$ (cm$^{-2}$)</th>
<th>$\log N(\text{Al} II)$ (cm$^{-2}$)</th>
<th>$\log N(\text{Si} II)$ (cm$^{-2}$)</th>
<th>$\log N(\text{Fe} II)$ (cm$^{-2}$)</th>
<th>Ref.$^a$</th>
</tr>
</thead>
<tbody>
<tr>
<td><strong>The Metal-Poor DLA Sample</strong></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>J0035–0918</td>
<td>2.34010</td>
<td>20.55 ± 0.10</td>
<td>15.47 ± 0.15</td>
<td>15.51 ± 0.06</td>
<td>14.96 ± 0.08</td>
<td>11.73 ± 0.05</td>
<td>13.41 ± 0.04</td>
<td>12.98 ± 0.07</td>
<td>2</td>
</tr>
<tr>
<td>J0311–1722</td>
<td>3.73400</td>
<td>20.30 ± 0.06</td>
<td>14.02 ± 0.08</td>
<td>$\leq 13.07$</td>
<td>14.70 ± 0.08</td>
<td>...</td>
<td>13.31 ± 0.07</td>
<td>$\leq 13.76$</td>
<td>1</td>
</tr>
<tr>
<td>J0831+3358</td>
<td>2.30364</td>
<td>20.25 ± 0.15</td>
<td>...</td>
<td>$\leq 12.78$</td>
<td>14.93 ± 0.05</td>
<td>12.19 ± 0.06</td>
<td>13.75 ± 0.04</td>
<td>13.33 ± 0.06</td>
<td>1,4</td>
</tr>
<tr>
<td>Q0913+072</td>
<td>2.61843</td>
<td>20.34 ± 0.04</td>
<td>13.98 ± 0.05</td>
<td>12.29 ± 0.12</td>
<td>14.63 ± 0.01</td>
<td>11.78 ± 0.03</td>
<td>13.30 ± 0.01</td>
<td>12.99 ± 0.01</td>
<td>3</td>
</tr>
<tr>
<td>J1001+0343</td>
<td>3.07841</td>
<td>20.21 ± 0.05</td>
<td>13.58 ± 0.02</td>
<td>$\leq 12.50$</td>
<td>14.25 ± 0.02</td>
<td>...</td>
<td>12.86 ± 0.01</td>
<td>12.50 ± 0.14</td>
<td>1</td>
</tr>
<tr>
<td>J1016+4040</td>
<td>2.81633</td>
<td>19.90 ± 0.11</td>
<td>13.66 ± 0.04</td>
<td>$\leq 12.76$</td>
<td>14.13 ± 0.03</td>
<td>...</td>
<td>12.90 ± 0.05</td>
<td>...</td>
<td>3</td>
</tr>
<tr>
<td>J1037+0139</td>
<td>2.70487</td>
<td>20.50 ± 0.08</td>
<td>...</td>
<td>13.27 ± 0.04</td>
<td>15.06 ± 0.04</td>
<td>12.32 ± 0.03</td>
<td>13.97 ± 0.03</td>
<td>13.53 ± 0.02</td>
<td>1</td>
</tr>
<tr>
<td>J1340+1106</td>
<td>2.50792</td>
<td>20.09 ± 0.05</td>
<td>...</td>
<td>12.83 ± 0.05</td>
<td>15.05 ± 0.03</td>
<td>12.29 ± 0.02</td>
<td>13.76 ± 0.02</td>
<td>13.51 ± 0.02</td>
<td>1</td>
</tr>
<tr>
<td>J1340+1106</td>
<td>2.79583</td>
<td>21.00 ± 0.06</td>
<td>...</td>
<td>14.08 ± 0.02</td>
<td>16.04 ± 0.04</td>
<td>13.27 ± 0.03</td>
<td>14.70 ± 0.02</td>
<td>14.32 ± 0.01</td>
<td>1</td>
</tr>
<tr>
<td>J1419+0829</td>
<td>3.04973</td>
<td>20.40 ± 0.03</td>
<td>...</td>
<td>13.28 ± 0.02</td>
<td>15.17 ± 0.02</td>
<td>...</td>
<td>13.83 ± 0.01</td>
<td>13.54 ± 0.03</td>
<td>1</td>
</tr>
<tr>
<td>J1558+4053</td>
<td>2.55332</td>
<td>20.30 ± 0.04</td>
<td>14.22 ± 0.06</td>
<td>12.66 ± 0.07</td>
<td>14.54 ± 0.04</td>
<td>11.92 ± 0.06</td>
<td>13.32 ± 0.02</td>
<td>13.07 ± 0.06</td>
<td>3</td>
</tr>
<tr>
<td>Q2206–199</td>
<td>2.07624</td>
<td>20.43 ± 0.04</td>
<td>14.41 ± 0.03</td>
<td>12.79 ± 0.05</td>
<td>15.05 ± 0.03</td>
<td>12.18 ± 0.01</td>
<td>13.65 ± 0.01</td>
<td>13.33 ± 0.01</td>
<td>3</td>
</tr>
<tr>
<td><strong>Literature DLAs</strong></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>Q0000–2620</td>
<td>3.39012</td>
<td>21.41 ± 0.08</td>
<td>...</td>
<td>14.70 ± 0.02</td>
<td>16.42 ± 0.10</td>
<td>...</td>
<td>15.06 ± 0.02</td>
<td>14.87 ± 0.03</td>
<td>5</td>
</tr>
<tr>
<td>Q0112–306</td>
<td>2.41844</td>
<td>20.50 ± 0.08</td>
<td>...</td>
<td>13.16 ± 0.04</td>
<td>14.95 ± 0.08</td>
<td>...</td>
<td>13.62 ± 0.02</td>
<td>13.33 ± 0.05</td>
<td>6</td>
</tr>
<tr>
<td>J0140–0839</td>
<td>3.69660</td>
<td>20.75 ± 0.15</td>
<td>14.13 ± 0.08</td>
<td>$\leq 12.38$</td>
<td>14.69 ± 0.01</td>
<td>11.82 ± 0.04</td>
<td>13.51 ± 0.09</td>
<td>12.77 ± 0.19$^b$</td>
<td>7</td>
</tr>
<tr>
<td>J0307–4945</td>
<td>4.46658</td>
<td>20.67 ± 0.09</td>
<td>...</td>
<td>13.57 ± 0.12</td>
<td>15.91 ± 0.17</td>
<td>13.36 ± 0.06</td>
<td>14.68 ± 0.07</td>
<td>14.21 ± 0.17</td>
<td>8</td>
</tr>
<tr>
<td>Q1108–077</td>
<td>3.60767</td>
<td>20.37 ± 0.07</td>
<td>...</td>
<td>$\leq 12.84$</td>
<td>15.37 ± 0.03</td>
<td>...</td>
<td>14.34 ± 0.02</td>
<td>13.88 ± 0.02</td>
<td>6</td>
</tr>
<tr>
<td>J1337+3153</td>
<td>3.16768</td>
<td>20.41 ± 0.15</td>
<td>13.98 ± 0.06</td>
<td>$\leq 12.80$</td>
<td>14.43 ± 0.09</td>
<td>12.00 ± 0.05</td>
<td>13.24 ± 0.05</td>
<td>13.14 ± 0.26</td>
<td>9</td>
</tr>
<tr>
<td>J1558–0031</td>
<td>2.70262</td>
<td>20.67 ± 0.05</td>
<td>...</td>
<td>14.46$^c$</td>
<td>15.86$^c$</td>
<td>...</td>
<td>14.24$^c$</td>
<td>14.11$^c$</td>
<td>10</td>
</tr>
<tr>
<td>Q1946+7658</td>
<td>2.84430</td>
<td>20.27 ± 0.06</td>
<td>...</td>
<td>12.59 ± 0.04</td>
<td>14.82 ± 0.01</td>
<td>...</td>
<td>13.60 ± 0.01</td>
<td>13.24 ± 0.01</td>
<td>11</td>
</tr>
<tr>
<td>Q2059–360</td>
<td>3.08293</td>
<td>20.98 ± 0.08</td>
<td>...</td>
<td>13.95 ± 0.02</td>
<td>16.09 ± 0.04</td>
<td>...</td>
<td>14.86 ± 0.05</td>
<td>14.48 ± 0.02</td>
<td>6</td>
</tr>
<tr>
<td>J2155+1358</td>
<td>4.21244</td>
<td>19.61 ± 0.10</td>
<td>13.95 ± 0.06</td>
<td>...</td>
<td>14.50 ± 0.05</td>
<td>11.92 ± 0.17</td>
<td>13.25 ± 0.04</td>
<td>12.93 ± 0.23</td>
<td>12</td>
</tr>
</tbody>
</table>

$^a$References as in Table 4.11.$^b$See comment in Table 4.11. $^c$An error estimate is not provided by the authors.
The [O/Fe] values for DLAs and stars in Figure 4.11 are listed in Table C.1 and C.2 respectively.

**Table C.1: [Fe/H] and [O/Fe] in VMP DLAs**

<table>
<thead>
<tr>
<th>QSO Name</th>
<th>$z_{abs}$</th>
<th>logN(H I)/cm$^{-2}$</th>
<th>[Fe/H]</th>
<th>[O/Fe]</th>
<th>Ref.$^a$</th>
</tr>
</thead>
<tbody>
<tr>
<td>Q0000−2620</td>
<td>3.39012</td>
<td>21.41 ± 0.08</td>
<td>−2.01 ± 0.09</td>
<td>+0.33 ± 0.10</td>
<td>2</td>
</tr>
<tr>
<td>J0035−0918</td>
<td>2.34010</td>
<td>20.55 ± 0.10</td>
<td>−3.04 ± 0.12</td>
<td>+0.76 ± 0.11</td>
<td>3</td>
</tr>
<tr>
<td>Q0112−306</td>
<td>2.41844</td>
<td>20.50 ± 0.08</td>
<td>−2.64 ± 0.09</td>
<td>+0.40 ± 0.09</td>
<td>4</td>
</tr>
<tr>
<td>J0140−0839</td>
<td>3.69660</td>
<td>20.75 ± 0.15</td>
<td>−3.45 ± 0.24</td>
<td>+0.70 ± 0.19</td>
<td>5</td>
</tr>
<tr>
<td>J0307−4945</td>
<td>4.46658</td>
<td>20.67 ± 0.09</td>
<td>−1.93 ± 0.19</td>
<td>+0.48 ± 0.24</td>
<td>6</td>
</tr>
<tr>
<td>J0311−1722</td>
<td>3.73400</td>
<td>20.30 ± 0.06</td>
<td>≤ −2.03</td>
<td>≥ −0.26</td>
<td>1</td>
</tr>
<tr>
<td>J0831+3358</td>
<td>2.30364</td>
<td>20.25 ± 0.15</td>
<td>−2.39 ± 0.16</td>
<td>+0.38 ± 0.08</td>
<td>1,7</td>
</tr>
<tr>
<td>Q0913+072</td>
<td>2.61843</td>
<td>20.34 ± 0.04</td>
<td>−2.82 ± 0.04</td>
<td>+0.42 ± 0.02</td>
<td>8</td>
</tr>
<tr>
<td>J1001+0343</td>
<td>3.07841</td>
<td>20.21 ± 0.05</td>
<td>−3.18 ± 0.15</td>
<td>+0.53 ± 0.14</td>
<td>1</td>
</tr>
<tr>
<td>J1037+0139</td>
<td>2.70487</td>
<td>20.50 ± 0.08</td>
<td>−2.44 ± 0.08</td>
<td>+0.31 ± 0.04</td>
<td>1</td>
</tr>
<tr>
<td>Q1108−077</td>
<td>3.60767</td>
<td>20.37 ± 0.07</td>
<td>−1.96 ± 0.07</td>
<td>+0.27 ± 0.04</td>
<td>4</td>
</tr>
<tr>
<td>J1337+3153</td>
<td>3.16768</td>
<td>20.41 ± 0.15</td>
<td>−2.74 ± 0.30</td>
<td>+0.07 ± 0.28</td>
<td>9</td>
</tr>
<tr>
<td>J1340+1106</td>
<td>2.50792</td>
<td>20.09 ± 0.05</td>
<td>−2.05 ± 0.05</td>
<td>+0.32 ± 0.04</td>
<td>1</td>
</tr>
<tr>
<td>J1340+1106</td>
<td>2.79583</td>
<td>21.00 ± 0.06</td>
<td>−2.15 ± 0.06</td>
<td>+0.50 ± 0.04</td>
<td>1</td>
</tr>
<tr>
<td>J1419+0829</td>
<td>3.04973</td>
<td>20.40 ± 0.03</td>
<td>−2.33 ± 0.04</td>
<td>+0.41 ± 0.04</td>
<td>1</td>
</tr>
<tr>
<td>J1558−0031</td>
<td>2.70262</td>
<td>20.67 ± 0.05</td>
<td>−2.03$^b$</td>
<td>+0.53$^b$</td>
<td>10</td>
</tr>
<tr>
<td>J1558+4053</td>
<td>2.55332</td>
<td>20.30 ± 0.04</td>
<td>−2.70 ± 0.07</td>
<td>+0.25 ± 0.07</td>
<td>8</td>
</tr>
<tr>
<td>Q1946+7658</td>
<td>2.84430</td>
<td>20.27 ± 0.06</td>
<td>−2.50 ± 0.06</td>
<td>+0.36 ± 0.02</td>
<td>11</td>
</tr>
<tr>
<td>Q2059−360</td>
<td>3.08293</td>
<td>20.98 ± 0.08</td>
<td>−1.97 ± 0.08</td>
<td>+0.39 ± 0.04</td>
<td>4</td>
</tr>
<tr>
<td>J2155+1358</td>
<td>4.21244</td>
<td>19.61 ± 0.10</td>
<td>−2.15 ± 0.25</td>
<td>+0.35 ± 0.24</td>
<td>12</td>
</tr>
<tr>
<td>Q2206−199</td>
<td>2.07624</td>
<td>20.43 ± 0.04</td>
<td>−2.57 ± 0.04</td>
<td>+0.50 ± 0.03</td>
<td>8</td>
</tr>
</tbody>
</table>

$^a$References as in Table 4.11. $^b$An error estimate is not provided by the authors.
<table>
<thead>
<tr>
<th>Star Name</th>
<th>$T_{\text{eff}}^a$</th>
<th>$\log g^b$</th>
<th>[Fe/H]</th>
<th>[O/Fe]</th>
<th>Ref. $^c$</th>
</tr>
</thead>
<tbody>
<tr>
<td>HD 2796</td>
<td>4950</td>
<td>1.50</td>
<td>$-2.37 \pm 0.10$</td>
<td>$+0.23 \pm 0.13$</td>
<td>1</td>
</tr>
<tr>
<td>HD 3567</td>
<td>6000</td>
<td>4.07</td>
<td>$-1.11 \pm 0.06$</td>
<td>$+0.29 \pm 0.06$</td>
<td>2</td>
</tr>
<tr>
<td>HD 4306</td>
<td>4990</td>
<td>3.04</td>
<td>$-2.24 \pm 0.10$</td>
<td>$+0.47 \pm 0.11$</td>
<td>3</td>
</tr>
<tr>
<td>HD 26169</td>
<td>4972</td>
<td>2.49</td>
<td>$-2.19 \pm 0.10$</td>
<td>$+0.34 \pm 0.09$</td>
<td>3</td>
</tr>
<tr>
<td>HD 27928</td>
<td>5044</td>
<td>2.67</td>
<td>$-2.05 \pm 0.10$</td>
<td>$+0.22 \pm 0.11$</td>
<td>3</td>
</tr>
<tr>
<td>HD 45282</td>
<td>5352</td>
<td>3.15</td>
<td>$-1.46 \pm 0.10$</td>
<td>$+0.33 \pm 0.07$</td>
<td>3</td>
</tr>
<tr>
<td>HD 97320</td>
<td>5976</td>
<td>4.16</td>
<td>$-1.16 \pm 0.06$</td>
<td>$+0.20 \pm 0.12$</td>
<td>2</td>
</tr>
<tr>
<td>HD 108317</td>
<td>5300</td>
<td>2.76</td>
<td>$-2.16 \pm 0.10$</td>
<td>$+0.49 \pm 0.12$</td>
<td>3</td>
</tr>
<tr>
<td>HD 111980</td>
<td>5694</td>
<td>3.99</td>
<td>$-1.03 \pm 0.06$</td>
<td>$+0.17 \pm 0.07$</td>
<td>2</td>
</tr>
<tr>
<td>HD 122563</td>
<td>4600</td>
<td>1.10</td>
<td>$-2.71 \pm 0.10$</td>
<td>$+0.31 \pm 0.13$</td>
<td>1</td>
</tr>
<tr>
<td>HD 126587</td>
<td>4712</td>
<td>1.66</td>
<td>$-2.76 \pm 0.10$</td>
<td>$+0.13 \pm 0.14$</td>
<td>3</td>
</tr>
<tr>
<td>HD 126681</td>
<td>5524</td>
<td>4.48</td>
<td>$-1.12 \pm 0.06$</td>
<td>$+0.33 \pm 0.08$</td>
<td>2</td>
</tr>
<tr>
<td>HD 128279</td>
<td>5336</td>
<td>2.95</td>
<td>$-2.10 \pm 0.10$</td>
<td>$+0.17 \pm 0.23$</td>
<td>3</td>
</tr>
<tr>
<td>HD 132475</td>
<td>5818</td>
<td>3.95</td>
<td>$-1.39 \pm 0.06$</td>
<td>$+0.33 \pm 0.09$</td>
<td>2</td>
</tr>
<tr>
<td>HD 140283</td>
<td>5690</td>
<td>3.69</td>
<td>$-2.32 \pm 0.06$</td>
<td>$+0.50 \pm 0.19$</td>
<td>2</td>
</tr>
<tr>
<td>HD 160617</td>
<td>5931</td>
<td>3.77</td>
<td>$-1.72 \pm 0.06$</td>
<td>$+0.22 \pm 0.15$</td>
<td>2</td>
</tr>
<tr>
<td>HD 166913</td>
<td>6039</td>
<td>4.11</td>
<td>$-1.50 \pm 0.06$</td>
<td>$+0.22 \pm 0.24$</td>
<td>2</td>
</tr>
<tr>
<td>HD 186478</td>
<td>4700</td>
<td>1.30</td>
<td>$-2.49 \pm 0.10$</td>
<td>$+0.47 \pm 0.10$</td>
<td>1</td>
</tr>
<tr>
<td>HD 189558</td>
<td>5613</td>
<td>3.91</td>
<td>$-1.07 \pm 0.06$</td>
<td>$+0.31 \pm 0.05$</td>
<td>2</td>
</tr>
<tr>
<td>HD 205650</td>
<td>5733</td>
<td>4.39</td>
<td>$-1.12 \pm 0.06$</td>
<td>$+0.30 \pm 0.09$</td>
<td>2</td>
</tr>
<tr>
<td>HD 213657</td>
<td>6114</td>
<td>3.85</td>
<td>$-1.86 \pm 0.06$</td>
<td>$+0.30 \pm 0.24$</td>
<td>2</td>
</tr>
<tr>
<td>HD 218857</td>
<td>5015</td>
<td>2.78</td>
<td>$-1.72 \pm 0.10$</td>
<td>$+0.18 \pm 0.09$</td>
<td>3</td>
</tr>
<tr>
<td>HD 274939</td>
<td>5090</td>
<td>2.79</td>
<td>$-1.43 \pm 0.10$</td>
<td>$+0.35 \pm 0.05$</td>
<td>3</td>
</tr>
<tr>
<td>HD 298986</td>
<td>6071</td>
<td>4.21</td>
<td>$-1.30 \pm 0.06$</td>
<td>$+0.23 \pm 0.17$</td>
<td>2</td>
</tr>
<tr>
<td>BD $-18^\circ 5550'$</td>
<td>4750</td>
<td>1.40</td>
<td>$-2.94 \pm 0.10$</td>
<td>$+0.08 \pm 0.24$</td>
<td>1</td>
</tr>
<tr>
<td>BD $-01^\circ 2582'$</td>
<td>5072</td>
<td>2.92</td>
<td>$-2.03 \pm 0.10$</td>
<td>$+0.31 \pm 0.11$</td>
<td>3</td>
</tr>
<tr>
<td>BD $+17^\circ 3248'$</td>
<td>5250</td>
<td>1.40</td>
<td>$-1.99 \pm 0.10$</td>
<td>$+0.46 \pm 0.11$</td>
<td>1</td>
</tr>
<tr>
<td>BD $+23^\circ 3130$</td>
<td>5170</td>
<td>3.00</td>
<td>$-2.29 \pm 0.06$</td>
<td>$+0.38 \pm 0.15$</td>
<td>2</td>
</tr>
<tr>
<td>CS 22186 - 035</td>
<td>4900</td>
<td>1.50</td>
<td>$-2.88 \pm 0.10$</td>
<td>$+0.26 \pm 0.26$</td>
<td>1</td>
</tr>
<tr>
<td>CS 22873 - 055</td>
<td>4550</td>
<td>0.70</td>
<td>$-2.87 \pm 0.10$</td>
<td>$+0.19 \pm 0.13$</td>
<td>1</td>
</tr>
<tr>
<td>CS 22891 - 209</td>
<td>4700</td>
<td>1.00</td>
<td>$-3.16 \pm 0.10$</td>
<td>$+0.41 \pm 0.16$</td>
<td>1</td>
</tr>
<tr>
<td>CS 22892 - 052</td>
<td>4850</td>
<td>1.60</td>
<td>$-2.91 \pm 0.10$</td>
<td>$+0.14 \pm 0.26$</td>
<td>1</td>
</tr>
<tr>
<td>CS 22896 - 154</td>
<td>5250</td>
<td>2.70</td>
<td>$-2.58 \pm 0.10$</td>
<td>$+0.64 \pm 0.23$</td>
<td>1</td>
</tr>
<tr>
<td>CS 22948 - 066</td>
<td>5100</td>
<td>1.80</td>
<td>$-3.01 \pm 0.10$</td>
<td>$+0.54 \pm 0.23$</td>
<td>1</td>
</tr>
<tr>
<td>CS 22949 - 037</td>
<td>4900</td>
<td>1.50</td>
<td>$-3.81 \pm 0.10$</td>
<td>$+1.54 \pm 0.13$</td>
<td>1</td>
</tr>
<tr>
<td>CS 22953 - 003</td>
<td>5100</td>
<td>2.30</td>
<td>$-2.73 \pm 0.10$</td>
<td>$+0.44 \pm 0.23$</td>
<td>1</td>
</tr>
<tr>
<td>CS 22966 - 057</td>
<td>5300</td>
<td>2.20</td>
<td>$-2.52 \pm 0.10$</td>
<td>$+0.70 \pm 0.20$</td>
<td>1</td>
</tr>
<tr>
<td>CS 22968 - 014</td>
<td>4850</td>
<td>1.70</td>
<td>$-3.42 \pm 0.10$</td>
<td>$+0.51 \pm 0.27$</td>
<td>1</td>
</tr>
<tr>
<td>CS 22941 - 053</td>
<td>4700</td>
<td>1.30</td>
<td>$-2.92 \pm 0.10$</td>
<td>$+0.43 \pm 0.17$</td>
<td>1</td>
</tr>
<tr>
<td>CS 22945 - 041</td>
<td>4800</td>
<td>1.50</td>
<td>$-2.71 \pm 0.10$</td>
<td>$+0.37 \pm 0.13$</td>
<td>1</td>
</tr>
<tr>
<td>CS 29516 - 024</td>
<td>4650</td>
<td>1.20</td>
<td>$-2.94 \pm 0.10$</td>
<td>$+0.28 \pm 0.20$</td>
<td>1</td>
</tr>
<tr>
<td>CS 29518 - 051</td>
<td>5200</td>
<td>2.60</td>
<td>$-2.58 \pm 0.10$</td>
<td>$+0.63 \pm 0.23$</td>
<td>1</td>
</tr>
<tr>
<td>CS 30325 - 094</td>
<td>4950</td>
<td>2.00</td>
<td>$-3.17 \pm 0.10$</td>
<td>$+0.36 \pm 0.33$</td>
<td>1</td>
</tr>
<tr>
<td>CS 31082 - 001</td>
<td>4825</td>
<td>1.50</td>
<td>$-2.79 \pm 0.10$</td>
<td>$+0.28 \pm 0.15$</td>
<td>1</td>
</tr>
</tbody>
</table>

$^a$Stellar effective temperature (K). $^b$Surface gravity (cgs).
$^c$References—1: Cayrel et al. (2004); 2: Nissen et al. (2002); 3: García Pérez et al. (2006).
References

Cooke, R., et al. 2010b, in prep
Dekker H., D’Odorico S., Kaufer A., Delabre B., Kotzlowski H., 2000, SPIE, 4008, 534
References

509, A93
References

RAS, 345, 480
Savage, B. D., & Sembach, K. R. 1996, ARAA, 34, 279
Schneider D. P., et al., 2007, AJ, 134, 102
Suda T., et al., 2008, PASJ, 60, 1159
Wiltshire D. L., 2009, PhRvD, 80, 123512