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ASTRONOMY
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# The chemical composition of the extreme halo stars

## I. Blue spectra of 20 dwarfs★

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**Summary.** The abundances of 13 elements in the atmospheres of 20 extreme halo dwarfs have been determined on the basis of spectra obtained with the ESO Cassegrain Echelle Spectrograph equipped with a CCD detector. The abundance trends for most elements are well defined and show a remarkably small scatter, which may be explained by the observational uncertainties alone. The main results are:

(1) The $\alpha$ elements Mg, Ca and Ti are overabundant with respect to Fe, by some $0.4-0.5$ dex.

(2) Al is overdeficient relative to Mg, by an amount which increases with increasing metal deficiency, at least down to $[\mathrm{Fe/H}]=-2.5$.

(3) The Cr/Fe ratio is solar at all metallicities.

(4) The s elements Sr, Y and Ba are overdeficient with respect to Fe in stars with $[\mathrm{Fe/H}]<-2.3$.

(5) When $[\mathrm{Fe/H}]>-2.3$, the s elements abundances relative to Fe are roughly constant at either solar or higher values. In particular, Zr is overabundant by about 0.5 dex.

(6) The $r$ process element Eu is overabundant with respect to Fe, by some 0.6 dex, but with a slightly larger scatter.

(7) The "nitrogen-rich metal-poor stars" show enhanced Al lines.

(8) The intercombination line of Mg I is affected by departures from LTE.

**Key words:** stellar abundances–Population II stars–chemical evolution of the Galaxy–non-LTE effects

### 1. Introduction

The determination of the chemical composition of the extreme halo stars is essential for the study of the early galactic evolution. It allows not only to put tight constraints on the theories of nucleosynthesis in metal-poor environment, but also to derive important information on the very first stellar generation(s) (Population III?) as well as on pre-stellar nucleosynthesis.

The advent of solid-state detectors, such as the Reticon and the CCD, have revolutionized the field of stellar spectroscopy, by allowing accurate data to be obtained for a large number of stars. As a consequence, our knowledge of the chemical evolution of the galactic disk has improved dramatically (see, e.g., Nissen et al., 1985). The situation is, however, much more obscure as far as the halo is concerned. This is due in large part to the apparent faintness of most halo stars. Two recent works have dealt with metal-poor stars, including halo objects. The study of François (1986a) includes 35 dwarfs, but only 9 of them belong to the halo (which we define as $[\mathrm{Fe/H}]<-1$), while only one star can be classified as "extreme" ($[\mathrm{Fe/H}]<-2$). On the other hand, Gratton and Sneden (1987, hereafter GS) analysed 46 stars, out of which 6 are intermediate halo dwarfs ($-2<[\mathrm{Fe/H}]<-1$). Rather arbitrarily, we consider as dwarfs the stars which have $\log g>3$; this definition includes some subgiants.

We are thus left with a fair number of intermediate halo dwarfs but with only one extreme metal-poor dwarf (namely, the well-known "subdwarf" HD 140283). To find an analysis including a significant number of extreme metal-poor dwarfs, we have to go back to Peterson (1981), whose sample contains 8 dwarfs with $[\mathrm{Fe/H}]<-2$. However, that study is based on much lower quality photographic or image tube spectra. Finally, the reanalysis of Magain (1987a) deals with 20 dwarfs, 12 of which have $[\mathrm{Fe/H}]<-2$, but is based on heterogeneous literature data. Figure 1 presents a graphical summary of the published analyses. Most studies not considered in this short summary include only one or two halo dwarfs.

This lack of a well analysed sample of extreme halo dwarfs has prompted us to undertake the present work. Our aim is to obtain accurate abundances for a representative sample of halo dwarfs and subgiants, the emphasis being put on the extremely metal-poor ones. Our sample was selected mainly on the basis of the Strömgren $m_1$ index indicating very low metallicity, while $b-y$ and $c_1$ were used to exclude giants and horizontal branch stars. The final sample includes 20 dwarfs and subgiants, 13 of which having, as far as we know, never been analysed in detail. Moreover, 12 of them have $[\mathrm{Fe/H}]<-2$. The basic observational data for the program stars are listed in Table 1.

We did not include any giant for the following reason: we want the derived surface composition to be as close as possible to the chemical composition of the gas out of which the stars were formed. This means that two requirements have to be met. First, the surface composition should not have been altered during the star's lifetime: in short and roughly, the closer to the main sequence the star is, the better it is. Secondly, the derived composition should be as close as possible to the actual surface composition. This means that the analysis should be as free as possible of systematic errors. Two important sources of such

★ Based on observations carried out at the European Southern Observatory (La Silla, Chile)

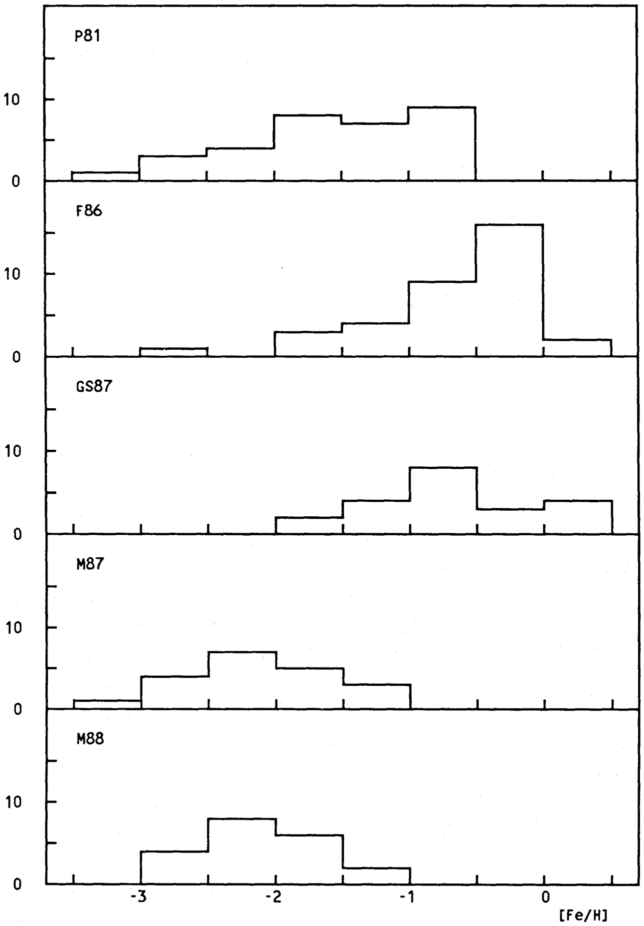


**Fig. 1.** Number of dwarfs analysed in some recent studies as a function of metallicity. P81 = Peterson (1981); F86 = François (1986a); GS87 = Gratton and Sneden (1987); M87 = Magain (1987a); M88 = this work

errors are: (1) possible departures from local thermodynamic equilibrium (LTE) and (2) errors in the adopted effective temperatures. As far as point (1) is concerned, all other parameters being equal, the lower pressures in giant atmospheres tend to favour departures from LTE. On the other hand, with regard to point (2), the smaller number of giants per unit volume in the Galaxy means that one has to go to larger distances to build a significant sample, thus increasing the reddening and, as a consequence, the uncertainty on the derived effective temperature. For these various reasons, we decided not to include any giant in our sample.

## 2. Observations and reductions

The observations were carried out with the Cassegrain Echelle Spectrograph (CASPEC) attached to the 3.6 m telescope at the European Southern Observatory (La Silla, Chile). The detector was a RCA CCD (type SID 501 EX, 320 × 512 pixels, 30 $\mu$m each). The 52 lines/mm Echelle grating was used in combination with the 300 lines/mm cross-disperser, thus providing a good separation between the orders. A 1.5 arcsec slit was used, giving a resolving power in excess of 15000 (FWHM) as measured on the stellar lines. The "blue" spectra discussed in this paper cover the range 3700–4700 Å. Spectra were also obtained in the range 5000–6000 Å but, due to the large amount of data, their analysis will be deferred to a subsequent paper. The S/N ratio is generally in excess of 100, except in the UV spectral region, where the exposure level is significantly lower.

The spectra were reduced with the help of the ESO, MIDAS and IHAP facilities, running on the VAX 750 and HP 1000 computers at La Silla. The standard MIDAS reduction package for Echelle spectra was slightly modified in order to yield optimum results in our particular case. Although wavelength calibration and flat field exposures had been obtained after each stellar exposure, they were not used in the final reduction. The wavelength calibration was performed on the stellar lines themselves, with the advantage that radial velocities are automatically corrected for. The internal precision of the calibration, as measured by the scatter around the fitting polynomial, was found to be nearly as good as the one based on the thorium lines, and perfectly adequate for the present analysis. Moreover, no flat-field was necessary for correcting pixel-to-pixel sensitivity variations, which are negligible in the blue spectral range on that particular CCD chip. On the other hand, flat-fielding was found to be inefficient for correcting the variation of exposure level due to the blaze of the Echelle grating. No standard star spectrum was obtained. No merging of the different orders into a single spectrum was attempted: all orders were treated independently, as single spectra. In summary, only the basic features of the MIDAS package were used: order location, background subtraction, order extraction, wavelength calibration. This rather basic reduction scheme was preferred because it allowed to keep the control on all the reduction steps and proved able to give quite accurate results, as this paper will demonstrate.

The subsequent reduction was performed with the help of the IHAP facility. By inspection of the solar atlases, a number (typically 20) of continuum windows were selected in each order. The mean exposure level was determined in each window and a low order Spline was fitted through these points. Note that for these extremely weak-lined stars, the continuum level is well determined, even in the blue spectral region. The division of the spectrum by the fitting curve yielded the normalised spectrum, on which the equivalent widths were measured by direct integration of the line profiles and by gaussian fitting in the case of lines with negligible damping wings ($W < 100$ mÅ). The final equivalent widths are weighted averages of these two measurements (with more weight on the second method). Sample spectra, for three stars with different metallicities, are presented in Fig. 2.

The measured equivalent widths, along with a detailed investigation of the quality for our data, will be presented in a separate paper. Let us just point out that a comparison with the equivalent widths of Magain (1985) for the two stars analysed in that paper shows a very good agreement, especially for the weak and medium lines ($W < 100$ mÅ), which are the best abundance indicators.

## 3. Determination of the model parameters

### 3.1. *Effective temperature*

The effective temperatures were determined from the $b-y$ and $V-K$ colour indices, using the new calibrations of Magain (1987b),

**Table 1.** Basic observational data

| Star | V | b-y | V-K | Obs. date | Exp. time (min) |
|---|---|---|---|---|---|
| HD 3567 | 9.25 | 0.330 | 1.35 | 13/14 Oct 86 | 20 |
| HD 16031 | 9.78 | 0.323 | 1.31 | 13/14 Oct 86 | 25 |
| HD 19445 | 8.05 | 0.351 | 1.39 | 13/14 Oct 86 | 14 |
| HD 34328 | 9.43 | 0.355 | 1.43 | 13/14 Oct 86 | 25 |
| HD 59392 | 9.72 | 0.337 | 1.34 | 09/10 May 86 | 35 |
| HD 74000 | 9.64 | 0.311 | 1.26 | 09/10 May 86 | 35 |
| HD 84937 | 8.30 | 0.300 | 1.21 | 09/10 May 86 | 12 |
| HD 116064 | 8.80 | 0.348 | 1.44 | 09/10 May 86 | 15 |
| HD 122196 | 8.74 | 0.354 | 1.43 | 09/10 May 86 | 12 |
| HD 140283 | 7.21 | 0.377 | 1.57 | 09/10 May 86 | 4 |
| HD 160617 | 8.74 | 0.344 | 1.39 | 09/10 May 86 | 13 |
| HD 166913 | 8.23 | 0.326 | 1.31 | 09/10 May 86 | 5 |
| HD 181743 | 9.69 | 0.346 | 1.37 | 13/14 Oct 86 | 30 |
| HD 194598 | 8.35 | 0.343 | 1.38 | 13/14 Oct 86 | 9 |
| HD 213657 | 9.67 | 0.317 | 1.28 | 09/10 May 86 | 23 |
| BD-10°0388 | 10.37 | 0.321 | 1.36 | 13/14 Oct 86 | 60 |
| BD+02°3375 | 9.95 | 0.354 | 1.46 | 09/10 May 86 | 35 |
| BD+03°0740 | 9.81 | 0.306 | 1.30 | 13/14 Oct 86 | 40 |
| BD+17°4708 | 9.47 | 0.325 | 1.40 | 13/14 Oct 86 | 25 |
| CD-33°3337 | 9.08 | 0.338 | 1.37 | 13/14 Oct 86 | 20 |

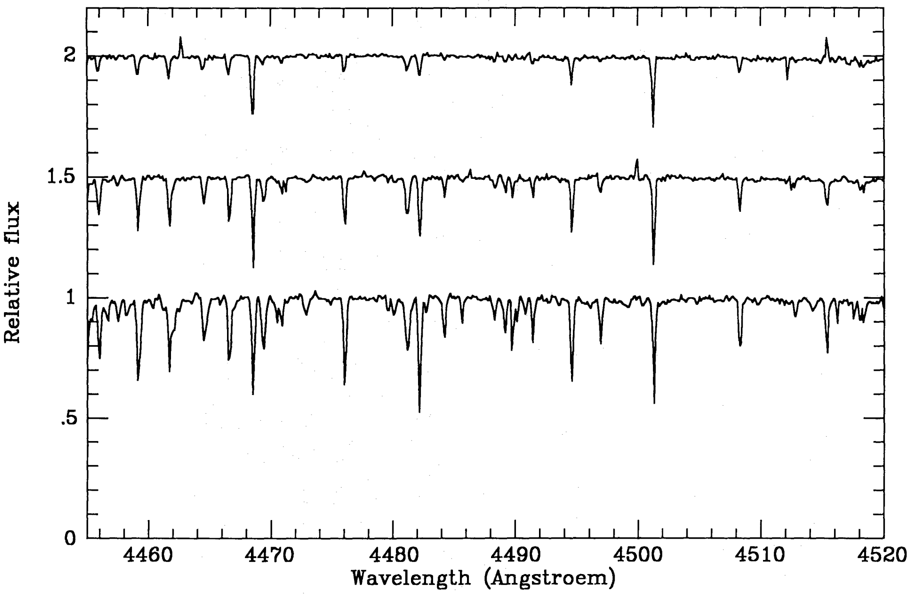


**Fig. 2.** Typical portion of the spectra of 3 stars with different metallicities. Upper spectrum: BD-10°388, [Fe/H] = −2.7, which is also the faintest of the sample. Middle spectrum: HD 16031, [Fe/H] = −2.0. Lower spectrum: HD 194598, [Fe/H] = −1.4

which are based on the infrared flux method (Blackwell and Shallis, 1977), and which are in close agreement with the calibrations of Carney (1983). The $b-y$ indices are from Carney (1983) of Lindgren (private communication). When available, the $V-K$ colours are also taken from Carney (1983). The stars not included in that paper have been observed either by the present author or by Th. Le Bertre with an InSb photometer at the ESO 1 m telescope. The agreement between the temperatures deduced from the two colour indices is quite good, the mean difference $\langle T(b-y) - T(V-K) \rangle$ amounting to 30 K, with an r.m.s. scatter of 60 K (on the individual values), which is thus a reasonable estimate of the random error affecting a single effective temperature determination.

As usual, zero reddening was assumed in the derivation of the effective temperature. This should be a reasonable approximation for these rather nearby dwarfs, although the temperature might be slightly underestimated in a few cases.

### 3.2. Surface gravity

Following the usual practice, the surface gravity was determined by forcing the lines of two ionization stages of the same element to indicate consistent abundances. Specifically, $\log g$ was changed until the Fe II lines gave the same abundance as the Fe I lines. The uncertainty on the resulting surface gravity, due to the scatter in the line abundances alone, is of the order of 0.2 to 0.3 dex. It was checked that the Ti ionization equilibrium gives essentially the same results, although with a slightly lower precision. It may be worth pointing out that these uncertainties are not due to the equivalent widths uncertainties alone, but that a non-negligible contribution comes from the oscillator strengths of the ionic lines.

Although there are some indications that departures from LTE may exist in the atmospheres of halo dwarfs (see, e.g., Magain, 1987c), we feel that, at the present stage, no other method is able to give gravities with any useful accuracy for extreme metal-poor stars. In particular, for most of the stars considered here, the parallaxes are very inaccurate or even unknown, while fitting the stars to any main sequence is completely arbitrary, especially around the turnoff, as the star might as well be a subgiant.

### 3.3. Metal abundance

The model overall metal abundance was determined from the present line data. The iron abundance was taken as representative of the mean metallicity and all abundances were scaled accordingly. Although of common use, this procedure is not completely justified since rather large variations occur in the relative abundances. However, this is of no consequence for the extreme metal-poor stars considered here since hydrogen is the main electron donor.

### 3.4. Microturbulent velocity

The microturbulent velocity was determined by forcing all the Fe I lines with Oxford $gf$ values (Blackwell et al., 1982, and references therein) and with $10\,\mathrm{m\AA} < W < 70\,\mathrm{m\AA}$ to give the same abundance. In order to avoid the systematic error discussed by Magain (1984c), the procedure recommended in that paper was adopted. Essentially the same result was obtained for all stars: the mean value for our sample amounts to $1.47\,\mathrm{km\,s^{-1}}$, with a scatter of $0.28\,\mathrm{km\,s^{-1}}$, which is comparable to the uncertainty affecting a single determination. In view of this fact, we adopt a value of $1.5\,\mathrm{km\,s^{-1}}$ for all stars. This is meant to reduce the errors due to a rather poorly determined value in a few stars (in which the Fe I lines are less than optimal for the microturbulent velocity determination), and has no serious consequence for the derived abundances. This mean value is in agreement with the determination of Magain (1985) for HD 19445 and HD 140283.

The adopted model parameters for the 20 program stars are listed in Table 2.

## 4. Method of analysis

### 4.1. Model atmosphere and line analysis

As in most recent analyses, the element abundance is obtained from each line by forcing the computed equivalent width to agree with the measured one. The usual assumptions of LTE and plane parallel (horizontally homogeneous) atmosphere are made. The models are interpolated in the grid of Magain (1983). These models were computed with a version of the MARCS program (Gustafsson et al., 1975) and are essentially identical with Gustafsson et al. models for metal-poor dwarfs. The interpolation procedure is the following. Only the $T(\tau)$ relation is interpolated in the grid. The other model quantities (electron and gas pressure, mean opacities) are computed from the temperature structure in

**Table 2.** Model parameters

| Star | $T_{\mathrm{eff}}$ | log g | [Fe/H] |
|---|---|---|---|
| HD 3567 | 5990 | 3.60 | -1.5 |
| HD 16031 | 6040 | 3.75 | -2.0 |
| HD 19445 | 5880 | 4.10 | -2.3 |
| HD 34328 | 5830 | 4.25 | -1.9 |
| HD 59392 | 5970 | 3.60 | -1.9 |
| HD 74000 | 6120 | 3.90 | -2.2 |
| HD 84937 | 6200 | 3.60 | -2.4 |
| HD 116064 | 5850 | 3.70 | -2.3 |
| HD 122196 | 5840 | 3.35 | -2.1 |
| HD 140283 | 5640 | 3.10 | -2.7 |
| HD 160617 | 5910 | 3.50 | -2.0 |
| HD 166913 | 6030 | 3.90 | -1.8 |
| HD 181743 | 5910 | 4.25 | -2.0 |
| HD 194598 | 5920 | 4.10 | -1.4 |
| HD 213657 | 6090 | 3.65 | -2.2 |
| BD-10°0388 | 6010 | 3.20 | -2.7 |
| BD+02°3375 | 5810 | 3.60 | -2.6 |
| BD+03°0740 | 6110 | 3.20 | -3.0 |
| BD+17°4708 | 5960 | 3.40 | -1.9 |
| CD-33°3337 | 5940 | 3.60 | -1.6 |

a consistent way, using the same routines as in the MARCS program. Whenever possible, the analysis is restricted to lines having equivalent widths between 10 and 70 mÅ. Using weaker lines would lead to increased random errors (and possibly some systematic overestimates), while stronger lines are very sensitive to microturbulence and damping. In a few cases, however, including Al, Sr and Ba, rather strong lines had to be used (except in the most metal-poor stars).

### 4.2. *Atomic parameters*

Our aim being to carry out as accurate as possible an analysis, relying on solar oscillator strengths is excluded for a number of reasons (see Magain, 1984a, 1985) which will be briefly summarized.

(1) Due to the extreme weakening of the lines in the program stars, the corresponding solar lines are too strong for an accurate oscillator strength to be determined since the latter would be very sensitive to errors in the equivalent widths, damping constants and microturbulent velocity.

(2) The blue part of the solar spectrum is so crowded that, in general, the local continuum cannot be reliably determined. In addition, some "missing opacity" (e.g., Holweger, 1970; Magain, 1983) may affect the derived oscillator strengths.

For these reasons, the use of solar oscillator strengths would introduce a large scatter in the derived abundances (see Magain, 1984a) and, very likely, some systematic errors. We are thus forced to rely on laboratory (or theoretical) oscillator strengths. Fortunately, many accurate *gf* values are now becoming available, so that the use of solar *gf* values is, in general, no more required.

In the very few cases where the oscillator strength has to be derived from the solar spectrum, the Holweger-Müller (1974, hereafter HM) model will be used, with a microturbulent velocity of 0.82 km s$^{-1}$, as derived from weak Fe I lines with Oxford *gf* values. The HM model is preferred to a theoretical solar model for the following reasons:

(1) we do not intend to make a differential analysis (which, let us stress it once more, would have no justification);

(2) what we want is a *gf* value which is as accurate as possible, and it has been shown in many occasions (see, e.g., Sauval et al., 1984) that the HM model is the one which gives the best results when used in conjunction with the LTE approximation.

However, we stress once more that we rely on solar *gf* values only when absolutely necessary and when, by chance, the solar line is not significantly blended and has a rather well defined local continuum. Further details on the oscillator strengths selected for the different elements are given in the appendix.

## 5. Estimation of the uncertainties

The sources of uncertainties can be divided in two categories. The first category includes the errors which act on a single line, like random errors in equivalent widths, oscillator strengths, damping constants, . . . , that is uncertainties on the line parameters. On the other hand, we put in the second category the errors which affect all the lines together, that is, mainly, the model errors (such as errors on the effective temperature, surface gravity, microturbulent velocity, overall metallicity, temperature structure, . . .).

Let us first discuss the line errors. For elements with first-class oscillator strengths and which are represented by a significant number of lines (mostly Fe I), the scatter of the deduced line abundances gives an estimate of the uncertainty coming from the random errors in the equivalent widths. Including all Fe I lines with 10 mÅ $< W <$ 70 mÅ, the scatter varies somewhat from star to star (from 0.06 to 0.11 dex) with a mean value of 0.08 dex. For unblended lines in this equivalent width range, the uncertainty on the derived abundance which is due to the random errors in the measured equivalent widths thus amounts to some 0.08 dex per line. The uncertainty on the element abundance is, of course, significantly smaller for elements with a large number of lines. This uncertainty is comparable to the values reported by François (1986a): 0.05 to 0.13 dex, and significantly smaller than the dispersion in Gratton and Sneden's (1987) abundances, which is of the order of 0.15 dex.

Despite the somewhat lower resolution of our spectra, and the additional complication in reducing multi-order, curved and tilted Echelle spectra, we thus achieve a precision which is at least as good as the one reported in these recent analyses. We believe that this is mainly due to (1) a very careful reduction procedure, (2) the rejection of all dubious lines and (3) the selection of the best atomic data available.

The uncertainties in the atomic data themselves are much more difficult to estimate. Random errors in oscillator strengths act exactly as random errors in equivalent widths and are included in the above estimate. Systematic errors are, of course, much more dangerous. When rather strong lines are used, additional uncertainties come from the (as usual poorly known!) damping constants. We consider, in particular, the atomic data for Ti II and Sc II to be far less than optimum.

Let us now turn to the second class of errors, i.e. due to model uncertainties. In view of the fact that the surface gravity is determined via the Fe lines, any error in one of the other parameters will have repercussions on the derived $\log g$, which have to be taken into account. In short, $\log g$ is not an independent parameter. The same is true for the overall metallicity, but on a negligible level, and will be neglected. Table 3 shows the effects on the derived abundances of changes of 100 K in effective temperature, 0.5 km s$^{-1}$ in microturbulent velocity, 0.3 dex in surface gravity and 0.5 dex in model metallicity in the case of HD 160617, which can be considered as a "mean star" in our sample. Also indicated is the combined effect of the likely uncertainties in all these quantities, i.e.: 60 K in $T_{\rm eff}$, 0.25 km s$^{-1}$ in $v_t$, 0.3 dex in $\log g$ and 0.25 dex in metallicity (the latter representing the effect of departures from scaled solar composition). These errors are supposed to be independent, with the restriction that for any error in another parameter, the effect on $\log g$ is taken into account. Table 3 shows that the resulting uncertainty on the relative abundances is never larger than 0.15 dex. Moreover, in all cases where it exceeds 0.1 dex, it happens that the abundance is deduced from rather strong resonance lines (Al, Sr, Ba).

Other sources of observational errors, such as continnum placement or background subtraction problems, should not play an important role in the present case. Moreover, the random part of these errors is included in the equivalent width uncertainties discussed at the beginning of this section.

Finally, there are some analysis uncertainties which are very difficult to quantify. In particular, the usual assumptions of LTE and horizontally homogeneous model atmosphere lack any real–physical–justification, except that we need them in order to reduce

**Table 3.** Influence of errors in the model parameters on the HD 160617 abundances

| Element ratio | $\delta T_{eff}$ =+100K | $\delta v_t$ =-0.5km/s | $\delta \log g$ =+0.3 | $\delta$[A/H] =+0.5 | overall uncertainty |
|---|---|---|---|---|---|
| [Fe/H] | +0.08 | +0.10 | 0.00 | +0.01 | 0.07 |
| [Mg/Fe] | -0.04 | -0.06 | -0.02 | 0.00 | 0.04 |
| [Al/Fe] | -0.02 | -0.02 | -0.10 | 0.00 | 0.10 |
| [Ca/Fe] | -0.02 | -0.03 | -0.01 | 0.00 | 0.02 |
| [Sc/Fe] | +0.03 | +0.13 | -0.02 | -0.02 | 0.07 |
| [Ti/Fe] I | +0.01 | -0.06 | 0.00 | +0.01 | 0.03 |
| [Ti/Fe] II | +0.02 | +0.06 | -0.01 | 0.00 | 0.03 |
| [Cr/Fe] | +0.01 | -0.03 | -0.01 | 0.00 | 0.02 |
| [Sr/Fe] | +0.02 | +0.08 | -0.13 | -0.02 | 0.14 |
| [Y/Fe] | +0.04 | +0.06 | -0.00 | +0.01 | 0.04 |
| [Ba/Fe] | +0.01 | +0.07 | -0.14 | -0.01 | 0.14 |
| [La/Fe] | +0.04 | -0.05 | -0.00 | +0.01 | 0.04 |
| [Eu/Fe] | +0.04 | -0.10 | 0.00 | +0.01 | 0.06 |
| [Al/Mg] | +0.02 | +0.08 | -0.08 | 0.00 | 0.09 |

Notes: (1) in the relative abundances, the same ionization stage is considered for both elements

(2) the overall uncertainty corresponds to the combined effect of likely errors in all parameters, i.e. 60K in $T_{eff}$, 0.25 in $v_t$, 0.3 in log g and 0.25 in [A/H] (the last one representing departures from scaled solar abundances).

a very complicated problem into a tractable one. Of course, this does not mean that they apply to real stars, and the fact that they are (nearly) always made should not hide their arbitrary character. Indeed, we have shown somewhere else (Magain, 1987c; see also Gustafsson, 1987) that significant departures from these assumptions might be present in halo stars' atmospheres. No attempt will be made to estimate the corresponding uncertainties (see, however, Sect.7), but it has to be reminded that, as in all similar analyses, the results are heavily dependent on the validity of these "classical" assumptions.

## 6. Results and discussion

The derived abundances are listed in Table 4. In this Sect., we will describe our main results and briefly discuss some of their implications for the chemical evolution of the Galaxy. Some points of vocabulary might not be out of place here. It is common practice to refer all abundances to the iron abundances, which is taken as a reference. "As deficient as Fe" means that the element abundance is reduced by the same factor as the Fe abundance is, with respect to the solar values. "Overabundant" (resp. "overdeficient") means less depleted (resp. more depleted) than Fe, unless explicitly specified (e.g., "overdeficient with respect to Mg").

### 6.1. *Iron abundances*

Due to the quality of the available oscillator strengths and the number of suitable lines, Fe is the element whose abundance should be known with the best accuracy. Our results for [Fe/H] are compared with some other recent determinations in Fig. 3. The mean difference for the 13 determinations in common amounts to −0.06 dex, with an r.m.s. scatter of 0.16 dex per individual datum (in the sense "this analysis minus other"). The analysis of HD 19445 by Gratton and Sneden (1987), although included in Fig. 3, was excluded from this mean since it was based on only a few lines and, according to the authors, should be interpreted with caution. In view of the facts that: (1) our estimate of the typical uncertainty in any of our [Fe/H] determinations is of the order of 0.08 dex and (2) the other analyses are, in general, of somewhat lower precision, the overall agreement is quite satisfactory. Within the framework of the LTE and classical model

**Table 4.** Element abundances

| Star | [Fe/H] | [Mg/Fe] | [Al/Fe] | [Ca/Fe] | [Sc/Fe] | [Ti/Fe] | [Cr/Fe] | [Sr/Fe] | [Y/Fe] | [Zr/Fe] | [Ba/Fe] | [La/Fe] | [Eu/Fe] |
|---|---|---|---|---|---|---|---|---|---|---|---|---|---|
| HD 3567 | -1.48 | +0.41 | -0.61 | +0.48 | +0.43 | +0.31 | +0.01 | +0.31 | -0.15 | +0.34 | +0.26 | +0.13 | +0.78 |
| HD 16031 | -1.97 | +0.52 | -0.59 | +0.51 | +0.31 | +0.53 | +0.04 | +0.34 | +0.04 | - | -0.02 | -0.21 | +0.72 |
| HD 19445 | -2.31 | +0.45 | -0.76 | +0.49 | +0.25 | +0.45 | -0.05 | +0.14 | - | - | -0.09 | +0.25 | +0.96 |
| HD 34328 | -1.86 | +0.42 | -0.43 | +0.31 | +0.34 | +0.37 | -0.05 | +0.26 | +0.16 | +0.53 | +0.10 | +0.12 | - |
| HD 59392 | -1.87 | +0.47 | -0.63 | +0.48 | +0.42 | +0.45 | +0.09 | +0.57 | +0.19 | +0.58 | +0.37 | +0.41 | +0.75 |
| HD 74000 | -2.23 | +0.46 | -0.27 | +0.45 | +0.31 | +0.43 | -0.04 | +0.49 | +0.13 | +0.68 | +0.07 | +0.25 | - |
| HD 84937 | -2.43 | +0.51 | -0.70 | +0.54 | +0.35 | +0.55 | +0.04 | +0.57 | 0.00 | - | -0.09 | +0.29 | +0.56 |
| HD 116064 | -2.28 | +0.59 | -0.71 | +0.55 | +0.33 | +0.50 | +0.08 | +0.26 | +0.01 | +0.47 | +0.02 | +0.15 | +0.86 |
| HD 122196 | -2.10 | +0.35 | -0.78 | +0.41 | +0.29 | +0.39 | +0.02 | +0.25 | -0.12 | - | -0.06 | -0.34 | - |
| HD 140283 | -2.75 | +0.51 | -0.85 | +0.34 | +0.25 | +0.25 | -0.20 | -0.34 | -0.58 | - | -0.96 | -0.05 | +0.21 |
| HD 160617 | -2.04 | +0.32 | -0.30 | +0.45 | +0.41 | +0.41 | -0.01 | +0.22 | +0.06 | +0.48 | +0.34 | +0.03 | +0.41 |
| HD 166913 | -1.80 | +0.50 | -0.32 | +0.46 | +0.49 | +0.43 | -0.01 | +0.42 | +0.21 | +0.58 | +0.22 | +0.11 | +0.49 |
| HD 181743 | -2.04 | +0.43 | -0.55 | +0.47 | +0.27 | +0.41 | +0.05 | +0.20 | - | +0.66 | -0.05 | +0.06 | - |
| HD 194598 | -1.37 | +0.27 | -0.48 | +0.31 | +0.07 | +0.24 | +0.05 | +0.07 | -0.03 | - | -0.07 | +0.02 | - |
| HD 213657 | -2.17 | +0.55 | -0.80 | +0.44 | +0.35 | +0.51 | +0.14 | +0.52 | -0.06 | - | 0.00 | -0.01 | +0.57 |
| BD-10°0388 | -2.66 | +0.52 | -0.72 | +0.60 | +0.45 | +0.55 | +0.04 | +0.33 | - | - | -0.21 | +0.27 | +0.85 |
| BD+02°3375 | -2.65 | +0.68 | -0.62 | +0.49 | +0.16 | +0.36 | +0.21 | +0.10 | -0.03 | - | -0.31 | +0.27 | - |
| BD+03°0740 | -2.98 | +0.54 | -0.83 | +0.48 | - | +0.40 | -0.18 | -0.25 | - | - | -0.67 | +0.65 | - |
| BD+17°4708 | -1.93 | +0.59 | -0.57 | +0.66 | +0.59 | +0.53 | +0.07 | +0.32 | -0.01 | +0.29 | +0.03 | -0.08 | +0.58 |
| CD-33°3337 | -1.63 | +0.50 | -0.40 | +0.44 | +0.41 | +0.38 | 0.00 | +0.34 | +0.07 | +0.44 | +0.18 | -0.12 | +0.58 |

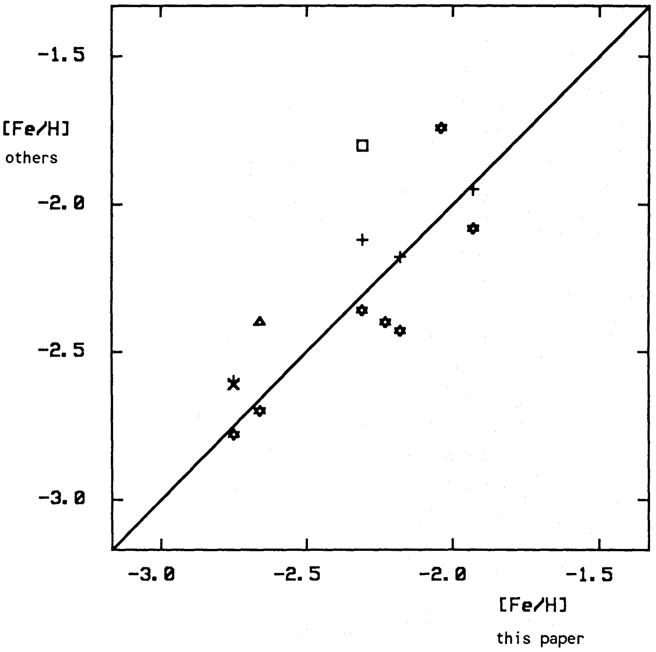


**Fig. 3.** Comparison of our results for [Fe/H] with those of other recent investigations. Peterson (1981): plusses; Barbuy et al. (1985): triangle; Francois (1986a): cross; Magain (1987a): stars; Gratton and Sneden (1987): square

atmospheres approximations, an abolute precision of 0.1 dex on [Fe/H] for extreme metal-poor stars is thus within reach of a careful analysis with modern techniques, even for 10th magnitude stars.

In a previous analysis based on photographic spectra and empirical model atmospheres (Magain, 1985), we obtained [Fe/H] $= -2.3$ for HD 19445 and $-3.0$ for HD 140283. While the agreement is excellent in the case of HD 19445, there is a 0.25 dex discrepancy for HD 140283. The latter may be, in large part, explained by the use of different models (the 1985 empirical model being roughly 200 K cooler than the present theoretical model in the line-forming layers). This discrepancy illustrates the rather large influence of the adopted model on the deduced absolute abundances [M/H]. Fortunately, the relative abundances [M/Fe] are much less affected (see Sects. 5 and 7 for more details).

### 6.2. *Alpha elements*

The abundances of Mg, Ca and Ti relative to Fe are plotted against the Fe abundance in Figs. 4 to 6, which show that these elements are clearly overabundant, in agreement with the results of most recent studies. We obtained the following mean overabundances:

$$[\mathrm{Mg/Fe}] = +0.47 \pm 0.09 \quad (1)$$

$$[\mathrm{Ca/Fe}] = +0.47 \pm 0.08 \quad (2)$$

$$[\mathrm{Ti/Fe}] = +0.40 \pm 0.09 \quad (3)$$

The quoted uncertainty is the r.m.s. scatter of the individual values (as will be throughout this paper, unless specified). It can be entirely accounted for by the observational and analysis uncertainties, so that there is no evidence for any cosmic variation or scatter in these abundances ratios throughout the range $-3.0 <$ [Fe/H] $< -1.5$.

These overabundances are in excellent agreement with the results of Magain (1987a), who found $+0.43 \pm 0.09$ for Mg and

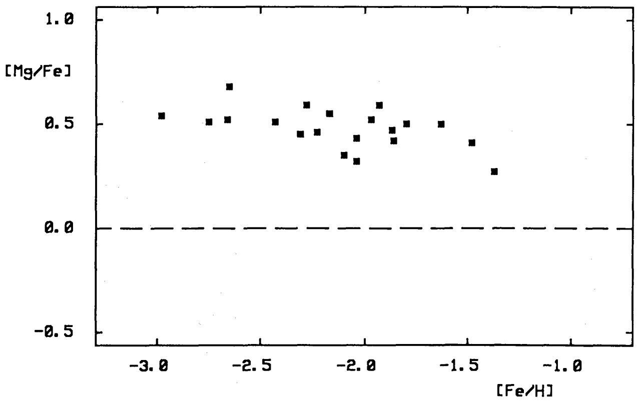


**Fig. 4.** [Mg/Fe] versus [Fe/H]

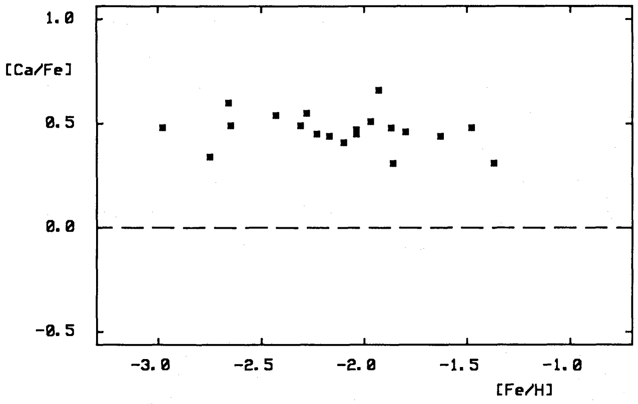


**Fig. 5.** [Ca/Fe] versus [Fe/H]

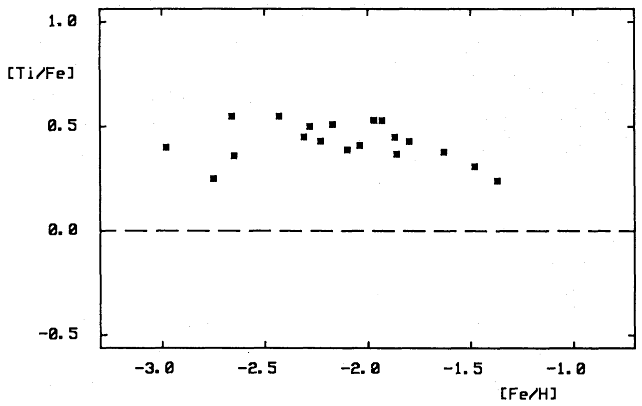


**Fig. 6.** [Ti/Fe] versus [Fe/H]

$+0.50 \pm 0.14$ for Ca. Our Mg overabundance is somewhat larger than the value found by François (1986a), according to whom $[Mg/Fe] = +0.35 \pm 0.10$ for 6 stars with $[Fe/H] < -1$. Gratton and Sneden (1987) also obtain lower overabundances ($+0.33 \pm 0.15$ for Mg, $+0.18 \pm 0.16$ for Ca and $+0.24 \pm 0.15$ for Ti). Note, however, that the latter results display a significantly larger scatter (the uncertainty listed in their Table 5 is the standard deviation of the mean and does not correspond to the scatter of the individual determinations). The reason for these discrepancies is not obvious. It is probably not due to differences in the observational material since the present results are in agreement with those of Magain (1987a) who used a completely different, and inhomogeneous, observational material. Two main differences distinguish our analyses from those of François and GS. First, the stars in our halo samples are generally more metal-poor than the ones studied by Francois and GS, so that the comparison is not completely meaningful. For example, in the analysis of GS, there seems to be a slight increase of the alpha element overabundances as the metallicity decreases (see, e.g., their Fig. 4). The second major difference is the fact that we make an absolute analysis (using laboratory or theoretical oscillator strengths) while they work differentially with respect to the sun. One could argue that our Mg oscillator strengths, coming from theoretical calculations, are not of the highest accuracy. However, Mg is just the element for which the discrepancy is smaller. On the other hand, for Ca, Ti and Fe, we use first class oscillator strengths from the Oxford group. So, it seems that part of the discrepancy might come from the use of a differential analysis, which is not suited to extreme metal-poor stars, as heavily dependent on any error in damping constants, solar continnum, equivalent widths and microturbulent velocity (not to consider possible departures from LTE, likelier in strong lines). Since we have discussed this point in much detail in previous papers (Magain, 1984a, 1985), we will not repeat the argument here.

### *6.3. Aluminium*

Although most authors agree on the fact that Al is overdeficient relative to Mg in metal-poor stars, there has been much debate in recent years on the question of whether this overdeficiency is constant or increasing with the Mg deficiency. The debate goes back to 1980 when Spite and Spite, analysing a sample of 6 metal-poor stars, found that $[Al/Mg] \sim 0$, in contrast to earlier Peterson's (1978) finding that [Al/Mg] was decreasing with overall metal abundance. A few years later, Arpigny and Magain (1983), in a detailed reanalysis of HD 19445 and HD 140283 (the first belonging to Peterson's sample and the second to the Spites') showed that the apparent discrepancy was due to a number of analysis problems. Their results were in qualitative agreement with Peterson's, although they found a less extreme overdeficiency. The increase of the Al overdeficiency with the Mg deficiency thus seemed well established.

All these analyses, however, used the (rather) strong UV resonance line(s) of Al I. Al abundances were later obtained by François (1986a) using red excited lines. Because these lines are weak, these data were restricted, however, to stars with $[Fe/H] > -2$. The conclusion of the author was that there was indeed an overdeficiency, but more or less constant at $[Al/Fe] \sim 0$, corresponding to $[Al/Mg] \sim -0.4$. This led to the suspicion that the resonance lines of Al might give systematically lower abundances than the excited lines. Such a trend has in fact been found for other elements (e.g., O, Mg, Ca, Fe) in the halo subgiant HD 76932 (Magain, 1987c).

However, it may be worth mentioning that the three most metal-poor stars analysed by Francois – and the only ones with $[Fe/H] < -1.4$, which are thus essential for defining the trend – are in fact binaries (Stryker et al., 1985). In our opinion, this makes the results for these stars somewhat doubtful: if both components of the binary system contribute to the flux, significant errors in the deduced abundance may be expected. Removing these three stars would mean that no trend can be deduced any more from François' data.

The results of GS, also based on excited lines, show a rather large scatter, and may be compatible with any of the solutions (constant or increasing overdeficiency). Finally, the reanalysis of Magain (1987a) showed that all previous data were compatible with a steady increase of the Al overdeficiency as the star becomes more metal-poor, but a possible discrepancy between resonance and excited lines could not be ruled out.

After this rather long introduction, let us now turn to the results of the present analysis. [Al/Mg] is plotted against [Mg/H] in Fig. 7. As usual when the 3961 Å resonance line is used, we obtain an Al overdeficiency increasing with decreasing [Mg/H]. Two stars, however, disagree with the average trend in having higher Al abundances. These stars are HD 74000 and HD 160617. Now comes the interesting point: these stars are among the 5 (or so) known "nitrogen-rich metal-poor stars" in the field, and the only ones in our sample. This naturally suggests that the N-rich stars are also Al-rich. Further confirmation of this hypothesis comes from François (1986b) who found another N-rich star to be Na-rich (Na and Al are believed to behave in the same way).

Some nitrogen enhancement has also been found in some globular cluster giants and a similar correlation with Na and Al has been observed in few cases (see, e.g., Smith, 1987, for a review of this problem). Several processes have been advocated to explain this phenomenon: internal mixing, mass transfer in close binaries or primordial N enhancement (see Laird, 1985; Carbon et al., 1987). The correlation with Al found here in the case of field dwarfs would seem to favour the last of these hypotheses.

Let us now come back to the variation of [Al/Mg] with [Mg/H]. Fig. 8 combines the present data with those of Magain (1987a). For the stars in common, the results of the present analysis are preferred in view of their higher precision and homogeneity. We stress that the results of François (1986a), as reanalysed by Magain (1987a), are included in this figure, at least for the stars with $T_{\mathrm{eff}} > 5000$ K and [Fe/H] $< -1$. The crosses indicate the N-rich stars while the open squares correspond to the known binaries. It is somewhat fascinating to find that *all* the stars falling significantly above the mean trend belong to one of these two categories. Removing them results in a nice trend of increasing Al overdeficiency with decreasing metallicity. We recall that Fig. 8 includes, as far as we know, all reliable [Al/Mg]

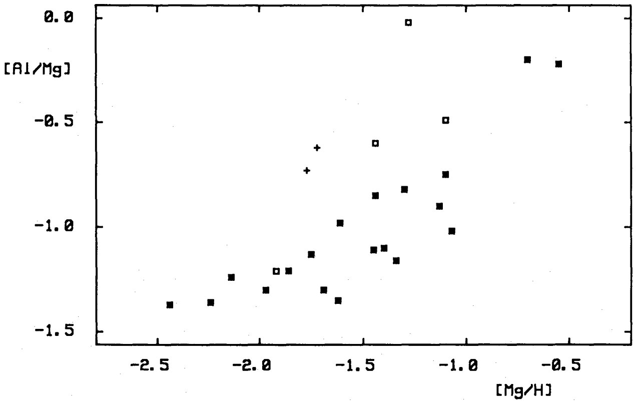


**Fig. 8.** Combination of the present data for [Al/Mg] versus [Mg/H] with those of Magain (1987a). Open squares indicate known binaries while plusses correspond to nitrogen-rich stars

determinations for dwarfs with [Fe/H] $< -1$ and $T_{\mathrm{eff}} > 5000$ K, either from the resonance line or from excited lines (although the last ones are few). Note that all dwarfs in GS sample for which [Fe/H] $< -1$ are either probable or certain binaries (Stryker et al., 1985) or included in the sample of François, whose data are of somewhat higher quality.

The presently available data thus indicate that, for single halo dwarfs without nitrogen enhancement, the Al overdeficiency increases with decreasing metallicity, in agreement with the predictions of explosive carbon burning computations (e.g., Arnett, 1971), at least down to [Mg/H] $\sim -2.0$, corresponding to [Fe/H] $\sim -2.5$. Below this value, our data seem to suggest some levelling off of the trend towards a constant value [Al/Mg] $\sim -1.3$. Such a constant value at very low metallicities is not incompatible with the theoretical expectations (Woosley: footnote in Bessell and Norris, 1984).

### *6.4. Scandium*

The results for Sc are displayed in Fig. 9, which shows that Sc is overabundant relative to Fe, with a mean value:

$$[\mathrm{Sc/Fe}] = +0.34 \pm 0.12 \tag{4}$$

A marginal odd-even effect seems to be present, the odd element Sc being slightly more deficient than the even elements Ca and

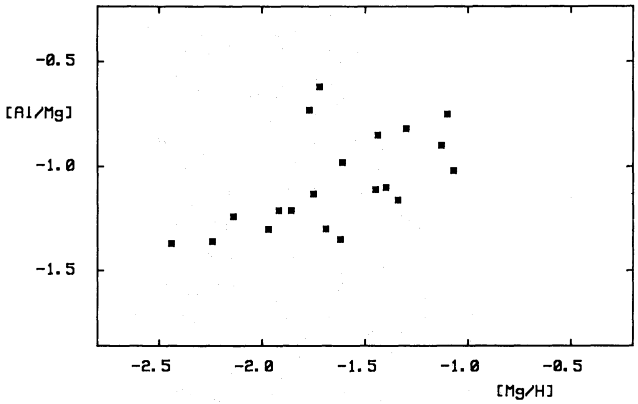


**Fig. 7.** [Al/Mg] versus [Mg/H]

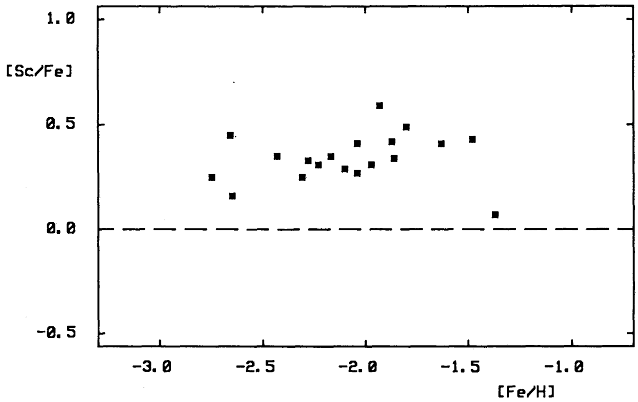


**Fig. 9.** [Sc/Fe] versus [Fe/H]

Ti, e.g.:

$$[\mathrm{Sc/Ca}] = -0.13 \pm 0.10 \qquad (5)$$

However, The Sc oscillator strengths are not of the highest quality, so that the Sc abundances should be considered with caution.

### 6.5. *Chromium*

Cr is the only iron peak element, besides Fe itself, for which we have obtained abundances. This element is thus important for assessing the representativity of Fe as an iron-peak element. The results are plotted in Fig. 10, which shows that the Cr/Fe ratio is essentially solar at all metallicities. The somewhat larger scatter at the low metallicity end may be explained by the fact that the number of Cr I lines available decreases with decreasing metallicity, a single line being left for the most metal-poor stars. The mean value is:

$$[\mathrm{Cr/Fe}] = +0.01 \pm 0.08 \qquad (6)$$

### 6.6. *The heavy elements*

The study of the elements synthesized by the *r* and *s* processes (rapid and slow neutron captures) was one of the main motivations for obtaining blue spectra of these extreme metal-poor stars. Indeed, most of their suitable lines fall in that spectral region.

In general, it is not possible to ascribe one of these elements to a single process. Most of them are represented by several isotopes, each isotope being, in many cases, produced by both processes, in variable proportions. The relative contributions of these two processes to the elemental abundances in solar system material have been estimated by Cameron (1982) and is summarized in Table 5 for the elements considered here. Although the relative contributions may be different in metal-poor gas, not very much is known on that subject, and we will tentatively adopt the classification suggested by Table 5, i.e., we will consider all these elements, except Eu, to be primarily *s*-process products.

In most past analyses, the heavy elements abundances were mainly obtained for giant stars (Spite and Spite, 1985). This may be explained by the fact that the lines available are often rather faint ion lines, which are more easily measurable in giant stars' spectra. The present study provides the first extensive set of heavy elements abundances for extreme metal-poor dwarfs. Since the atmospheric composition of the giants is more likely to have been

**Table 5.** Fractional contributions of the *s* and *r* processes to the solar system abundances (after Cameron, 1982)

| Element | s process | r process |
|---|---|---|
| Sr | 0.92 | 0.08 |
| Y | 0.73 | 0.27 |
| Zr | 0.79 | 0.21 |
| Ba | 0.88 | 0.12 |
| La | 0.65 | 0.35 |
| Eu | 0.09 | 0.91 |

modified as a result of stellar evolution (see Gustafsson, 1987), our data allow more reliable comparisons to the models of galactic chemical evolution. As such, they may provide important clues concerning the origin of these elements and the nature of the first stellar generations. Our results are presented in Figs. 11 to 16 and will now be discussed element by element.

Due to the strength of the available lines and the uncertainties affecting the *gf* values as well as the damping constants, our results for Sr should be considered as preliminary (Fig. 11).

It is generally assumed, following previous investigations (e.g., Barbuy et al., 1985) that the Y/Fe ratio is solar down to

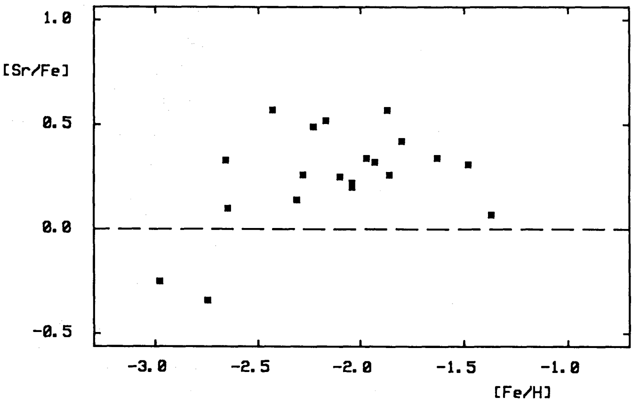


**Fig. 11.** [Sr/Fe] versus [Fe/H]

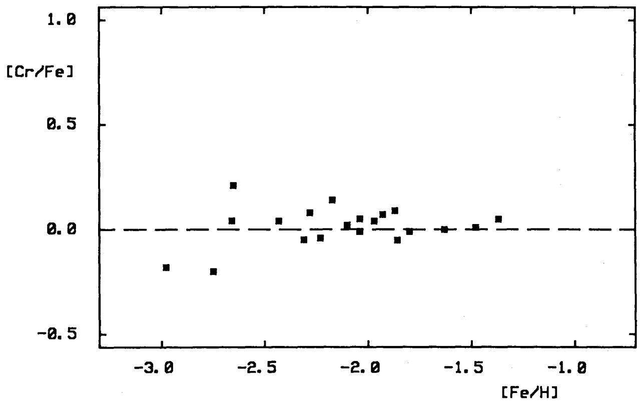


**Fig. 10.** [Cr/Fe] versus [Fe/H]

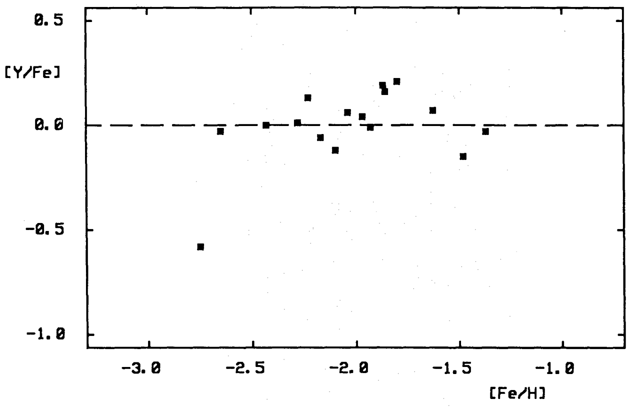


**Fig. 12.** [Y/Fe] versus [Fe/H]

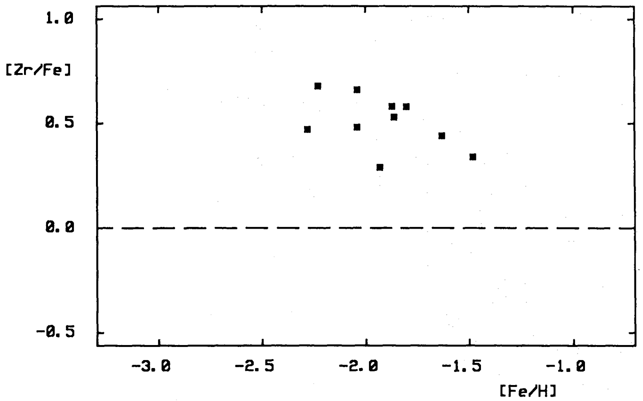


**Fig. 13.** [Zr/Fe] versus [Fe/H]

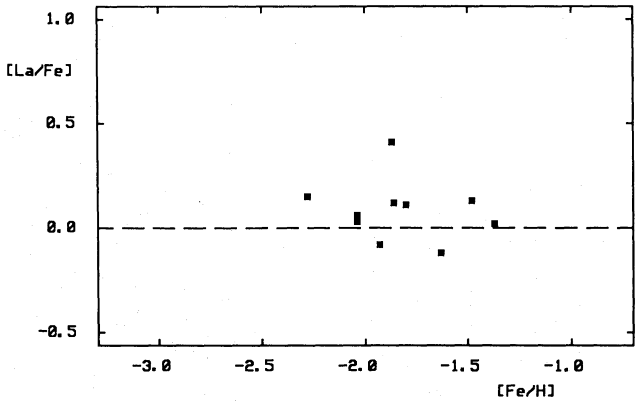


**Fig. 15.** [La/Fe] versus [Fe/H]

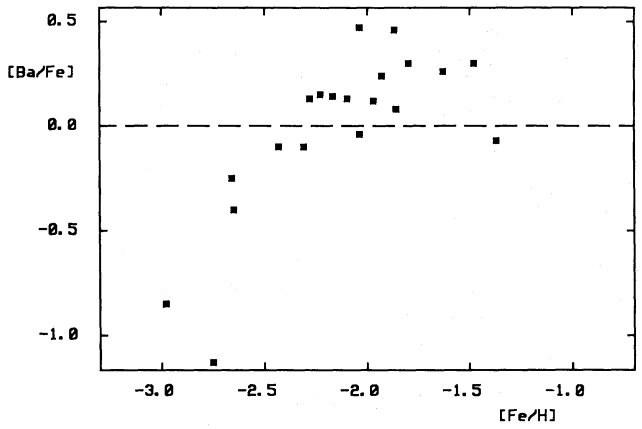


**Fig. 14.** [Ba/Fe] versus [Fe/H]

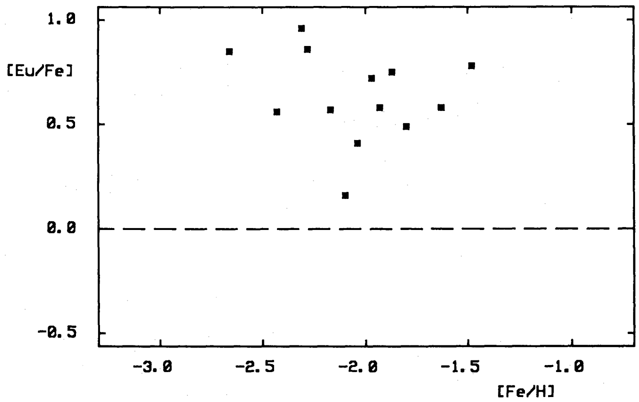


**Fig. 16.** [Eu/Fe] versus [Fe/H]

[Fe/H] $\sim -1.5$ and, then, starts to decrease with a slope of the order of 0.4. Our data show a slightly, but significantly, different behaviour (Fig. 12): Y/Fe is solar down to [Fe/H] $\sim -2.5$. At lower metallicities, yttrium is probably overdeficient. The only star showing a significantly negative [Y/Fe] is HD 140283. The other stars with similar metallicities are hotter and their Y lines are too weak to be measured with acceptable accuracy, they are thus missing from Fig. 12. However, the upper limits on unmeasurable lines confirm the reality of the Y overdeficiency for [Fe/H] $\lesssim -2.5$. As far as the less extreme stars are concerned, the mean value is (excluding HD 140283):

$$[\mathrm{Y/Fe}] = +0.05 \pm 0.11 \qquad (7)$$

As usual, the quoted uncertainty is the r.m.s. scatter of the individual values. Thus, despite the fact that most Y lines are located in the UV region, the scatter is only slightly larger than in the case of lighter elements.

Zirconium is clearly overabundant in halo stars (Fig. 13):

$$[\mathrm{Zr/Fe}] = +0.49 \pm 0.12 \qquad (8)$$

It may be worth noting that the Zr abundance is determined from a single weak Zr II line ($\lambda$ 4209), which is not measurable in our spectra as soon as [Fe/H] $\lesssim -2.3$. Although our results are based on a single line, several facts argue in favour of their reliability: (1) this line is not significantly blended and the local continuum is reasonably well defined; (2) in the sun, this line gives a Zr abundance which is in agreement with the mean value (Biémont et al., 1981): this confirms that it is not blended; (3) the small scatter in our data confirms (if necessary) that we are not measuring noise features (and after all, the line is not that weak). As far as we know, this is the first time that such a large Zr overabundance in halo dwarfs is clearly established. Its interpretation does not seem obvious. In particular, one might wonder if the similarity with the alpha elements overabundance is no more than a coincidence.

Thanks to the investigations of Spite and Spite (1978, 1985; see also Barbuy et al., 1985), the behaviour of Ba relative to Fe is rather well documented: Ba/Fe is essentially solar down to [Fe/H] $\sim -1.5$ where Ba starts to become overdeficient, the slope of the [Ba/Fe] versus [Fe/H] relation being of the order of 1.0. Our results (Fig. 14) are in qualitative agreement with these trends, with two significant differences. First, the overdeficiency starts at [Fe/H] $\sim -2.2$ instead of $-1.5$. Secondly, Ba seems to be slightly overabundant in the metallicity domain $-2.0 <$ [Fe/H] $< -1.5$. Fitting a straight line by least squares through the points with [Fe/H] $< -2$ leads to the following result:

$$[\mathrm{Ba/Fe}] = 1.29\,([\mathrm{Fe/H}] + 2.26) \qquad ([\mathrm{Fe/H}] < -2) \qquad (9)$$

while the mean value for the stars with [Fe/H] $> -2$ amounts to:

$$[\mathrm{Ba/Fe}] = +0.21 \pm 0.16 \qquad ([\mathrm{Fe/H}] > -2) \qquad (10)$$

with no significant slope. Note, however, that in this last metallicity domain, the Ba abundance is particularly sensitive to the adopted damping constant.

The difference between our results and those of Spite and Spite (1978) and others is mainly due to a different choice of the atomic parameters (oscillator strength and damping constant), different method of analysis (absolute versus differential) and a difference in the [Fe/H] values (see Sect. 4.2 and Magain, 1984b, 1985). The barium abundance in halo stars will be investigated in much more detail in a subsequent paper (Magain, 1988, in preparation) on the basis of very high quality data for weak excited lines. At the present stage, the best one can do is to conclude that the determination of the Ba abundance from the resonance line is affected by a rather large uncertainty, which is reflected by the difference between the results of different authors.

The lanthanum abundance (only measurable for stars with $[\mathrm{Fe/H}] \gtrsim -2.3$) is roughly solar and independent of metallicity (Fig. 15):

$$[\mathrm{La/Fe}] = +0.08 \pm 0.14 \tag{11}$$

As for Zr, the abundance is based on a single line and cannot be determined for the stars which show the Ba overdeficiency. The results for La are thus compatible with those for Ba.

The relative abundance of the *r* process element europium is plotted in Fig. 16, which shows the largest overabundance obtained in this study, but with the largest scatter:

$$[\mathrm{Eu/Fe}] = +0.64 \pm 0.19 \tag{12}$$

In particular, the scatter is larger than in the case of Zr and La, whose abundances are also determined from a single line, with a similar equivalent width. The very large hyperfine structure of the Eu line might account for some of this additional scatter, although a real cosmic scatter may not be ruled out. Other investigators (e.g. Luck and Bond, 1985) have also reported an unusually large scatter in [Eu/Fe].

The interpretation of the heavy elements abundances in terms of the existing models is far from obvious. Naive models for the *s* elements abundances predict that the points in the [*s*/Fe] versus [Fe/H] diagram should follow a straight line of slope 1 passing through the solar point $[s/\mathrm{Fe}] = 0$ at $[\mathrm{Fe/H}] = 0$. This behaviour basically reflects the secondary character of these *s* elements, i.e. the fact that they are produced in proportion to the pre-existing iron-peak elements. The observations are in complete disagreement with this simple model.

Another problem is the need for two prior stellar generations in order to account for the presence of secondary elements in the atmospheres of the stars considered: a first stellar generation is needed to produce iron-peak elements from primordial matter, while a second generation, containing these iron-peak elements, would be able to synthesize the *s* elements, which would thus appear in the atmospheres of the dwarfs of the third generation. Where are the stars of these first two generations?

Some recent models (e.g. Jones, 1985; Cayrel, 1986) give an elegant explanation for the lack of stars of the supposed first generation, i.e., stars with no metals. The massive stars would form more rapidly than the low mass stars and would complete their evolutionary cycle before these last ones are formed. The massive stars would then explode and pollute the interstellar gas out of which the first generation low mass stars would condense. This accounts for the absence of low mass stars with no metals. However, these models do not easily explain the presence of secondary elements in the atmospheres of these stars, except if a second generation of massive stars could again form out of the polluted gas, evolve and explode before any low mass star could form.

Another solution of this problem was proposed by Truran (1981) according to whom we do not observe, in the extreme metal-poor stars, any product of the *s* process, but rather the contribution of the *r* process (which is minor in solar system matter but would dominate in Pop II stars). Some detailed determinations of relative abundances in individual Pop II giants (Sneden and Parthasarathy, 1983; Sneden and Pilachowski, 1985) seem to confirm the predominance of the *r* process contribution. The abundances in individual stars of our sample can also be fitted by an *r* process distribution with a smaller *s* process contribution. For example, around $[\mathrm{Fe/H}] \sim -2$, a reasonable fit is obtained by scaling Cameron's solar abundances (Table 5) down by 1.3 dex for the *r* process and 1.9 dex for the *s* process.

However, if that hypothesis can reasonably explain the distribution of the relative abundances in individual stars, it completely fails to reproduce the variation of these relative abundances with [Fe/H]. As we already pointed out earlier (Magain, 1985), if the (supposedly primary) *r* process dominates in Pop II stars and if the (secondary) *s* process contribution increases with increasing metallicity, one would expect a constant overdeficiency of the heavy elements in the most extreme metal-poor stars followed by a gradual rise towards unit slope as the metallicity increases. This is not far from being exactly the reverse of the observed behaviour for Sr, Y and Ba. Moreover (and more generally), any variation in the relative abundances of these extreme metal-poor stars is difficult to explain in the framework of a model assuming that they all belong to the same (first) generation.

A much better fit to the *s* elements trends is provided by Twarog's (1981) model, which assumes that the so-called *s* elements are indeed produced by the *s* process (i.e. are secondary), but that their production relative to Fe is much less efficient in the disk than in the halo. This model is able to explain the unit slope in the most extreme Pop II stars followed by a more or less constant value (more precisely, it was built for). It also predicts some overabundance of the *s* process elements in the transition region, which seems to be confirmed by our data for Sr and Ba. Our results would imply that the production efficiency changed when [Fe/H] reached a value of the order of $-2$, which is not at the transition between the halo and the disk, but this is probably a rather minor modification. However, if Twarog's model is able to reproduce nicely the observed trends, it is completely ad hoc: this change in the efficiency of the *s* process at some value of the metallicity needs to be explained by nucleosynthesis calculations. Moreover, this model does not alleviate the need for two prior stellar generations.

## 7. Departures from LTE

As already pointed out, our results (as well as the results of all similar analyses) are heavily dependent on the validity of the assumptions of local thermodynamic equilibrium (LTE) and horizontally homogeneous model atmosphere (see Gustafsson, 1987). Possible departures from these assumptions were reported in our earlier work on HD 19445 and HD 140283 (Magain, 1984a, 1985) where it was shown that they could be interpreted as model errors, although departures from LTE could not be ruled out. Much more convincing effects were found in the star of intermediate metal deficiency HD 76932 (Magain, 1987c) where clear variations of the deduced abundances with the line excitation potential were established.

Our present data also allow some limited tests of these assumptions, by checking the consistency of the element abundances as derived from different sets of lines.

Let us first examine the case of the Fe I lines. Two different effects were reported in our previous works. In our analysis of the extreme halo stars HD 19445 and HD 140283, a systematic decrease of the Fe abundance with increasing excitation potential was found in the case of rather low excitation Fe I lines ($\chi < 2.6$ eV), with a slope of the order of $-0.035$ dex/eV. Our present results are compatible with such an effect, but its significance is only marginal: the weighted mean slope for the 20 stars considered is $-0.023$ dex/eV, with a scatter of 0.023 dex/eV. Although the empirical models used in our previous work (Magain, 1985) allow to correct for this effect, we decided not to use them here for a number of reasons: (1) they are ad hoc (although not more than some assumptions used in the computation of the theoretical models), (2) the changes in the relative abundances due to the use of these models instead of the theoretical ones are practically negligible, (3) the use of models which differ somewhat from the ones used by other investigators may be misleading, and one might think that our results would not be directly comparable to the ones of the other investigators.

The second effect, which was reported by Magain (1987c), is an increase of the deduced abundance with excitation potential for Fe I lines with $\chi > 2.6$ eV. Our present data do not include enough high excitation lines to confirm this effect.

Much better evidence for departures from LTE is found in the case of the Mg I lines. It has already been reported (e.g., Barbuy et al., 1985) that "the abundance computed from the intercombination line at 4571 Å, seems systematically and significantly lower [than the abundance computed from the excited lines]". The present data allow to investigate this effect in more detail. The abundance difference $\delta$[Mg/H] (= Mg abundance deduced from the excited lines minus abundance deduced from the intercombination line) is plotted as a function of the star's surface gravity in Fig. 17. Similar plots as a function of effective temperature or metallicity show no obvious trend. Figure 17 shows a clear correlation, the abundance difference being much larger in the stars with lower surface gravity. A linear least-squares fit gives the following results:

$$\delta[\mathrm{Mg/H}] = 1.81 - 0.41 \log g \qquad (13)$$

with an r.m.s. scatter of 0.10 dex, which can be accounted for by the observational errors only. The slope is significant at the 5 $\sigma$ level. This correlation is a very strong indication that departures from LTE are responsible for the difference in abundances. Indeed, as the surface gravity increases, so does the gas pressure, which means that the relative efficiency of the collisional processes with respect to the radiative processes increases and tends to establish LTE. We can thus safely conclude that either the intercombination line, or the excited lines, or both groups, are affected by departures from LTE. Since the comparison with the other elements abundances shows no dependence on $\log g$ when the excited lines are used, and does when the intercombination line is considered, it is very likely that this is the last one which is affected by departures from LTE.

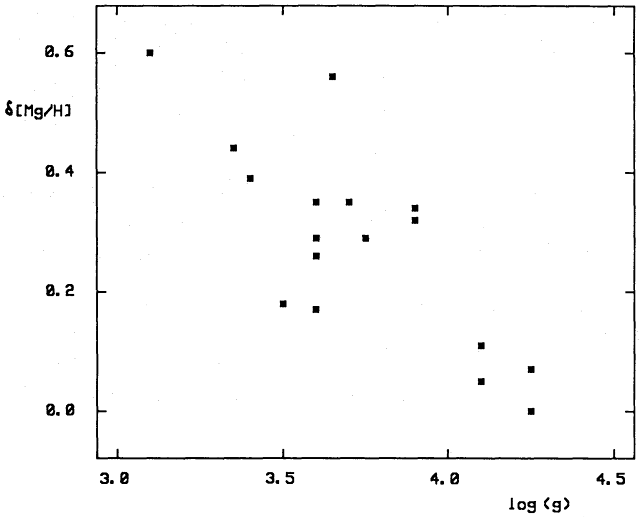


**Fig. 17.** Difference in the Mg abundance as given by the excited lines and by the intercombination line as a function of the stellar surface gravity

## 8. Conclusions

The main abundance trends in extreme Pop II dwarfs ([Fe/H] $< -1.5$) may be summarized in the following way.

(1) The alpha elements Mg, Ca and Ti are overabundant with respect to Fe. This overabundance is constant at about 0.4 to 0.5 dex, throughout the range $-3.0 <$ [Fe/H] $< -1.5$, with an extremely small (if any) cosmic scatter.

(2) For single dwarfs without nitrogen enhancement, Al is overdeficient relative to Mg. This overdeficiency increases with decreasing metallicity, in agreement with the predictions of explosive carbon burning, at least down to [Fe/H] $\sim -2.5$. Below this value, [Al/Mg] seems to tend towards a constant value of the order of $-1.3$.

(3) The Cr/Fe ratio is solar at all metallicities.

(4) For [Fe/H] $\gtrsim -2.3$, the abundances of the $s$ elements relative to Fe are roughly constant at either solar or higher values, Zr showing a particularly significant overabundance.

(5) For [Fe/H] $\lesssim -2.3$, the abundances of those $s$ elements which can be measured (Sr, Y, Ba) decrease faster than the Fe abundance, with a slope $d[s/\mathrm{Fe}]/d[\mathrm{Fe/H}]$ of the order of 1.

(6) The $r$ process element Eu is overabundant with respect to Fe, by some 0.6 dex. This overabundance seems independent of metallicity but the scatter is larger than for the other elements.

Despite many years of efforts and of interactions between theory and observation, no model seems able to account satisfactorily for all the observed variations in relative abundances, especially in the case of the most extreme halo stars. In particular, it is not clear whether the halo stars belong to an homogeneous population (either a single generation or several generations with a continuous evolution in chemical composition) or should be divided into two distinct populations. In this context, it is interesting to note that the behaviour of several elements seems to change at some rather well defined value of [Fe/H] $\sim -2.3$ (e.g. Al, Sr, Y, Ba and possibly some others, see Lambert, 1986; Gustafsson, 1987). Does this value of [Fe/H] represent a transition between two distinct populations?

On the other hand, it is worth stressing once more that all these results are dependent on the validity of the assumptions adopted in the model atmosphere analyses. In particular, the assumption of LTE needs to be checked carefully. Departures from LTE have been clearly established in some cases (i.e. the intercombination line of Mg I discussed in Sect. 7). The consequences of these departures from LTE on the deduced abundances should be investigated in more detail: the abundance trends not only have to be explained, they first need to be firmly established!

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## Appendix. Details on the individual elements

### 1. Light elements

The magnesium abundance is derived from excited (4.3 eV) Mg I lines, with oscillator strengths from Froese-Fisher (1975) and a damping width $\gamma_6$ chosen equal to 1.5 times the Unsöld value (i.e. a "damping enhancement factor" $f_6 = 1.5$, which corresponds to a damping constant $C_6$ equal to 2.8 times the Unsöld value, see Gray, 1976, p. 239). The Mg I intercombination line at 4571 Å was not used for the abundance determination as it is affected by departures from LTE (Sect. 7). Several determinations of its $gf$ value are available but show large discrepancies. Fortunately, the solar line is not too strong, neither badly blended and the local continuum is reasonably well defined. With a solar Mg abundance of 7.62 in the usual logarithmic scale with an hydrogen abundance $\log(A_{\rm H}) = 12.00$ (Lambert and Luck, 1978), we derive $\log(gf) = -5.43$, which is close to the value selected by Wiese et al. (1969): $\log(gf) = -5.40$. Moreover, we adopt $f_6 = 1.5$, as for the excited lines.

The only lines available for determining the aluminium abundance are the two violet resonance lines of Al I. The 3944 Å line has been shown by Arpigny and Magain (1983) to be blended by CH lines and, thus, unsuitable for abundance determination. We thus use the single 3961 Å line, with $\log(gf) = -0.34$ (Wiese and Martin, 1980) and a damping enhancement factor $f_6 = 1.6$, determined from a fit of the solar line, using a solar abundance of 6.49 (Lambert and Luck, 1978) and taking into account the absorption due to the nearby H line of Ca II.

The calcium abundance is deduced from excited lines of Ca I, with accurate oscillator strengths from Smith and O'Neill (1975) and Smith and Raggett (1981). The damping enhancement factors are taken as $f_6 = 1.5$ for the 1.9 eV lines and 1.8 for the 2.5 eV lines (Smith and Raggett, 1981). The solar abundance, computed from weak Ca I lines with oscillator strengths from the same sources, amounts to 6.36, in agreement with Smith (1981).

The oscillator strengths of the Sc II lines are from the NBS compilations (Wiese and Martin, 1980; Wiese and Fuhr, 1975). We adopt $f_6 = 1.5$ and a solar abundance of 2.99 (Holweger, 1979).

The titanium abundance is deduced from weak Ti I lines with accurate oscillator strengths from the Oxford group (Blackwell et al., 1986a, and references therein) and $f_6 = 1.5$. The solar abundance, as deduced from weak Ti I lines with Oxford oscillator strengths, amounts to 5.05. A number of Ti II lines are present in our spectra, but the lack of accurate oscillator strengths precludes a reliable abundance to be obtained for that ion. The $gf$ values are adopted from the critical compilation of Magain (1985). Only in the case of the most metal-poor star, BD +3°740, which has no Ti I line measurable, is the Ti abundance deduced from the Ti II lines.

### 2. Iron peak elements

The oscillator strengths of the Cr I lines are from the Oxford group (Blackwell et al., 1986b, and references therein), $f_6$ is 1.5 and the solar abundance, determined from weak lines with the same $gf$ values, amounts to 5.69.

The iron abundance is determined from Fe I lines with $10\,\text{mÅ} < W < 70\,\text{mÅ}$. The oscillator strengths of the low excitation lines ($\chi < 2.6$ eV) are from the Oxford group (Blackwell et al., 1982, and references therein). For the higher excitation lines ($\chi > 2.8$ eV), we use the oscillator strengths of Bridges and Kornblith (1974), corrected according to the formula given by Blackwell et al. (1982). The damping enhancement factor was determined for the low excitation lines from fits of the solar lines of the same excitation, and found to vary between 1.0 and 1.3, similarly to the results of Simmons and Blackwell (1982). For the high excitation lines, we adopt $f_6 = 1.5$ or 2.0 according to the parity of the lower level (even or odd). The solar abundance, as determined from weak Fe I lines with Oxford $gf$ values, is 7.68. The oscillator strengths of the Fe II lines are taken from Moity (1983), with a correction of 0.07 dex (Magain, 1985), determined from the lifetime measurements of Hannaford and Lowe (1982).

### 3. Heavy elements

For the Sr II lines, we adopt the oscillator strengths of Wiese and Martin (1980), $f_6 = 1.5$ and a solar abundance of 2.93 (Holweger, 1979).

The oscillator strengths and solar abundances for Y II and Zr II are from Hannaford et al. (1982) and Biémont et al. (1981).

The only line available for the determination of the barium abundance is the rather strong resonance line of Ba II. The oscillator strength is taken from Wiese and Martin (1980), the damping enhancement factor ($f_6 = 3.0$) from Holweger and Müller (1974) and the solar abundance (2.18) from Holweger (1979). The hyperfine structure was taken into account following Biehl (1976).

The oscillator strength of the only La II line measurable on our spectra ($\lambda$ 4086.7) was deduced from the solar line, with $f_6 = 1.5$ and a solar abundance of 1.08 (Holweger, 1979).

The europium abundance is determined from the 4129.7 Å resonance line, with the oscillator strength of Biémont et al. (1982) and $f_6 = 1.5$. The important hyperfine structure was approximately taken into account by increasing the microturbulent velocity from 1.5 to 8.0 km s$^{-1}$, as deduced from the width of the solar line. A solar abundance of 0.51 was obtained by Biémont et al. (1982) and is adopted.

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