HR 710 (HD 15144): An ultra-Sr-rich, magnetic Ap star with a close companion

C. R. Cowley\(^1\), † S. Hubrig\(^2\)

\(^1\)Department of Astronomy, University of Michigan, Ann Arbor, MI 48109-1042, USA
\(^2\)European Southern Observatory, Casilla 19001, Santiago 19, Chile

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ABSTRACT

The magnetic, chemically peculiar A-star HR 710 (HD 15144, ADS 1849A, AB Cet) has a close companion in a nearly circular orbit. Its 3-day period is unusually short for such stars. The system emits moderately hard x-rays, which likely come from a white dwarf secondary (ADS 1849Ab). The Sr\(^{ii}\) spectrum is very strong, and the resonance lines show similar core-nib structure to the stronger Ca K line. We place only loose constraints on a model. There are indications of a lower electron/gas pressure than expected from the star’s parallax and brightness. Strong-line profiles and anomalous excitation/ionization indicate significant deviations from traditional models, fixed by \(T_{\text{eff}}\), \(\log g\), and abundances. Weak spectral-line profiles and wavelength shifts probably indicate abundance patches rather than the presence of a secondary (ADS 1849Ab) spectrum. Our UVES spectra are from only one epoch. Additional low-noise, high-resolution spectra are needed. We discuss the spectrum within the context of abundance stratification.

Key words: stars: atmospheres–stars: chemically peculiar–stars: magnetic fields – stars: individual: HR710

1 INTRODUCTION

Roman (1949) listed HR 710 (HD 15144) among the probable members of the Ursa Major Stream. She noted the strengths of \(\lambda\lambda 4077\) and 4215. More recently, King, et al. (2003) enter a question mark in the column for “Final Memb.” for HD 15144A. Tokovinin (1997) concludes that the system ADS 1849A is a physical double, rather than a common proper motion pair, but the period must be very long. The primary, ADS 1849A (or itself a spectroscopic binary (Aa and Ab) with a well-determined period of 2.998 days (Bonsack 1981, Tokovinin 1997). This a very short period for a magnetic Ap star. One system, HD 200405 (cf. Carrier, et al. 2002), with an even shorter period is known–1.64 days. The HR 710 system appears to be single lined, so there has been a large uncertainty in the mass and luminosity of the companion. Hünsch, Schmitt, and Voges (1998, henceforth HSV) list HR 710 in their ROSAT survey of main sequence and subgiant stars of the Bright Star Catalogue. They found an x-ray luminosity of \(4.2 \times 10^{29}\) erg s\(^{-1}\). This is among the top third of the group. The hardness ratio they report, 0.12, is in the upper 15%. Thus HR 710 is significant, though not outstanding among these x-ray emitters.

Czesla and Schmitt (2007) report Chandra observations of HR 710 as well as the companion ADS 1849B in the energy band 0.5 to 2 keV Their x-ray flux for the primary is in reasonable agreement with the ROSAT data.

Extensive magnetic field measurements were made by Babcock (1958), and Bonsack (1981). The field varied between \(-0.45\) and \(-1.10\) kG—it did not reverse. Bonsack’s plot of \(H_{\text{eff}}\) folded on the 3-day orbital period appears random. These results are consistent with nearly magnetic and rotational pole-on observations. Bonsack sought other periods, and found marginal evidence for a 15.88-day period. He concluded \(< H_{\text{eff}} > \approx -700\) G, with an rms error of a single observation of 140 G.

One of us, CRC, measured wavelengths of a 2.4 Å mm\(^{-1}\) spectrum from the Dominion Astrophysical Observatory in 1979. Contemporary notes include a remarks on the strengths of Sr\(^{ii}\) and Ca\(^{ii}\) K. Additionally, they note the subordinate Sr\(^{ii}\) line, 4305.45 “has huge wings,” (see Fig. 1). Our description is similar to that of Babcock (1958), who also pointed out that Lines of Sr\(^{ii}\), like the K line, have strong, sharp cores with wide shallow wings. Much of the present study follows this lead.

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† E-mail: cowley@umich.edu (CC); shubrig@eso.org (SH)
strong subordinate lines? If so, would the study of these features add to our understanding of the structure of the photospheres of magnetic stars? The sections below deal with these questions.

We can place only loose constraints on the physical conditions in the atmosphere of HR 710. In addition to line core/wing and ionization and excitation anomalies, there is evidence for abundance patches on the surface, which can be seen in some of the line profiles.

2 OBSERVATIONS

The ESO spectra used in this study were obtained on 19 September 2005 with the UV-Visual Echelle Spectrograph (UVES) at UT2 of the VLT. The spectra used the new (November 2004) standard Dichroic settings were described by CHCGW, covering 3290 to 9460 Å. This configuration makes it possible to measure the λ8542-line, which had previously been unavailable because of an order gap. The resolving power is 80 000 in the blue-violet, and 110 000 in the red. The S/N ratios of the spectra differ, but range from about 100 to 400.

3 REDUCTIONS

The UVES spectra were reduced by the UVES pipeline Data Reduction Software (version 2.5), which is an evolved version of the ECHELLE context of MIDAS. The spectra were mildly Fourier filtered before the wavelengths were measured at Michigan.

We choose the continuum by picking high points of the relative flux using console displays 40 Å wide. In some cases, a Paschen line or water-vapor band would take up most of the display, and no continuum point would be chosen. The high points were connected by a spline and by linear interpolation. After some experimentation the linear fit was given weight 2, the spline 1, and the weighted average adopted. The continuum height is surely uncertain to a few per cent. Nevertheless we note that in the illustrations that follow, no additional (ad hoc) adjustments were made to the continua.

Wavelength measurements from 4527 to 4652 Å are from 2.4 Å mm^{-1} spectra (12766 U, and V) obtained at the Dominion Astrophysical Observatory in 1979. The instrument has been described by Fletcher et al. (1980).

4 THE ATMOSPHERE OF HR 710

4.1 Preliminary remarks

Ap stars were once thought to have abnormal abundances, but normal atmospheres. We now know that this is not the case. The profiles of the low Balmer and the strong metallic lines cannot be reproduced by standard plane-parallel atmospheres. Simple adjustments to temperature, gravity, and composition do not allow one to resolve inconsistencies in ionization and excitation. These difficulties have become apparent relatively recently. Horizontal chemical inhomogeneities (abundance patches) have been known in magnetic stars for many decades.

Extensive observations by Babcock (1958) and Bonsack...
We used several codes based on Strömgren or Geneva photometry to try to fix \( T_{\text{eff}} \) and \( \log g \). In addition, we considered ionization balance, Balmer, and Paschen profiles. The various implementations give a fairly wide range of values for \( T_{\text{eff}} \) (8400 to 8600 K) and \( \log g \) (3.0 to 4.0).

An implementation of Geneva Photometry by Kaiser (2006), based on Künzli, et al. (1997) includes an adjustment for \([\text{Fe}/\text{H}]\). The code gives values of \( \log g \) that decrease with increasing \([\text{Fe}/\text{H}]\). Allowed entries for \([\text{Fe}/\text{H}]\) are \(-1, 0, \text{and} +1\). The relation for both \( \log g \) and \( T_{\text{eff}} \) is very nearly linear. We used photometric measurements for HR 710 from the Simbad data base to obtain second-order fits, with very small quadratic terms. With \([\text{Fe}/\text{H}] = +1.0\), we obtain \( T_{\text{eff}} = 8460 \) K and \( \log g = 3.6 \). This somewhat low gravity is supported by our ionization balance of neutral and singly ionized Fe, where we get best agreement with \( \log g = 3.0 \). This ionization balance (see below) gives a temperature nearly 200 K higher.

Codes based on Strömgren photometry give significantly higher gravities, though they do not, to our knowledge, include a provision to allow for metallicity.

The higher surface gravities agree with Hubrig, et al. (2007), who give \( \log g = 4.2 \). The temperature estimate from this study is 8433 K.

### 4.3 Balmer lines

Hydrogen line strengths depend sensitively on temperature in cooler stars, and on surface gravity in hotter ones. HR 710 is at the high-temperature end of the region where the line strengths change from primarily temperature to primarily gravity sensitivity.

All Balmer calculations are based on the profiles from Stehlé & Hutcheon (1999).

Fig. 2 shows the observed and calculated profile of H\(\alpha\) for models with \( T_{\text{eff}} = 8600 \) K, and \( \log g = 3.0 \) and 4.0. The fit is arguably better with the lower-gravity model (upper smooth curve). Metallic lines have been omitted from the calculation to avoid confusion. The tendency of the lower Balmer cores of cooler magnetic Ap stars to be deeper, and sharper than the observations was called the core-wing anomaly (Cowley, et al. 2001). LTE calculations of the H\(\alpha\) core are never as deep as the observations. In the case of HR 710, departures of the calculations from the observed profiles in the regions outside the inner core are of interest. For H\(\alpha\) and especially for H\(\gamma\) and H\(\delta\) the calculations do not fit the observations until the far wings are reached.

The case for the lower gravity is somewhat stronger with H\(\gamma\) (Fig. 3) and H\(\delta\) (not shown).

Fig. 4 shows the region of the Balmer confluence. H18 (3691.56) is surely present, and H19 (3686.83) possibly so, though the feature is clearly strengthened by blending primarily with Cr \(\Pi\) 3687.03 and Fe I-75 3687.10. The Multiplet Number (75) is from Moore (1945). The strong Cr \(\Pi\) line does not appear in the Multiplet Tables. The Inglis-Teller relation (cf. Cowley 1970, Eq. AI-5) yields \( N_e = 13.4 \) if the last Balmer line is H19. This electron density is reached in our models with \( T_{\text{eff}} = 8600 \) K, and with \( \log g = 4 \) at a logarithmic optical depth \( \log \tau_{3640} = -0.7 \). For the \( \log g = 3 \) model the corresponding optical depth is \(-0.08\).

We have used the optical depth at \( \lambda 3640 \), short of the Balmer Jump. This should closely approximate the opacity where the confluence occurs, including the opacity due to overlapping higher Balmer lines. The basis of this statement may be found, among other places, in Cowley (2000, see also Appendix A).

We expect the electron density obtained by the Inglis-Teller formula would correspond to \( N_e \) at optical depth unity for the wavelength of the Balmer jump. Thus, the last countable Balmer line appears to give mild support for a lower-log \( g \) model. However a plot similar to Fig. 5 for \( \gamma \) Equ shows Balmer components possibly as high as H20. Moreover, in two other main-sequence stars, \( \nu \) Cap (B9.5 V) and \( \sigma \) Boo (F2 V), we can count to H22 or H23 and H19 or 20 respectively. A detailed study of the usefulness of the last countable Balmer line as an indicator of surface gravity would be worth while.
4.4 Paschen lines

The lower Paschen members are beyond our wavelength coverage. The first usable line is $\lambda 9229$, P9. The lowest Paschen members, P9, P10, and P11, are clearly fit better with a log $g = 4$ model than log $g = 3$. All Paschen line calculations used Lemke’s (1997) revised VCS profiles (Vidal, Cooper & Smith 1973), because the Stéhely–Hutcheon tables do not extend to large enough values of $N_e$ for the higher Paschen lines.

At $\lambda 8750$, P12, the lower-gravity model provides a better fit to the wings than the log $g = 4$ model, but the core is too deep. If we were to lower the observed continuum by 2 to 3%, the overall fit, core and wings is significantly better with the higher-gravity model. Fig. 4 shows the Paschen line at $\lambda 8598$, P14, which falls between the longer-wavelength pair of the IRT. Again, the higher-gravity model provides a better fit. One may surely count Paschen lines to P18. It is arguable whether higher Paschen members are present. Thus the limiting Balmer and Paschen series are in agreement with one another, but a definitive interpretation awaits further study. Better observational material, specifically taken and reduced for the purpose of investigation of series confluence would also be useful.

The Paschen lines P13, P14, P15, and P16 are blended with present special problems because they are near the series confluence. Details are explained in Appendix A.

5 IONIZATION EQUILIBRIUM

Lines of Fe I and II were chosen from a compilation of oscillator strengths by Fuhr and Wiese (2006). We used our line identification list (see URL below) to exclude cases with obvious blends. We then calculated abundances for models with log $g = 3$ and 4 for $T_{\text{eff}}$ ranging from 8000 to 8800 K. Calculations for all of these cases were made with, and without stratification. The stratification assumed was based on an analytic formula:

$$g(x) = b + (1 - b) \left[ \frac{1}{2} \sqrt{\pi/a} + \frac{1}{2} \text{erf} \left( \sqrt{a \cdot |x + d|^2} \right) \right].$$ (1)

In any given layer, the abundance of an element is multiplied by this $g(x)$. Here, $x = \log (T_{500})$, and the negative sign is taken for $x \leq -d$. This function is a smoothed step of height $1 - b$, with the center of the jump located at depth $\log T_{5000} = -d$. At large depths, $g(x)$ is nearly unity, while for the smallest depths, it has the value $b$. The Gaussian 1/e width of the step is $1/\sqrt{a}$.

For all spectra other than those of Ca II and Sr II, “standard parameters” were used: $a = 6.7$, log $b = -4$, and $d = 1$. We cannot claim any special validity for these particular values; they were used in CHCGW to reconcile profiles of the Ca K-line and one line of the IRT. We have already explored a wide domain of parameter space, and must leave possible adjustments of the stratification parameters to account for anomalous ionization and excitation to future work.

Tables of wavelengths and equivalent widths may be found on our web page:

http://www.astro.lsa.umich.edu/~cowley/hr710

For iron, abundances were found for 5 temperatures and 2 gravities, above and below the Balmer Jump, and with and without stratification using the standard parameters. We made plots of abundances versus equivalent widths and excitation potential, and used the former plots to fix the microturbulent parameter. Trends for iron were removed by the value $\xi = 3.5$. This value is somewhat large, but it also must account for the Zeeman broadening, which was not large enough to yield resolved patterns.

The results are illustrated in Fig. 5. Here, and elsewhere in this paper, ‘$N_{\text{tot}}$’ means the sum of relative abundances of all elements. If log(H) = 12, then $N_{\text{tot}} = 1.03 \times 10^{12}$. We used 70 lines of Fe I and 32 lines of Fe II, all with wavelengths longer than 3650 Å. This plot is the primary basis for consideration of low-gravity (or low pressure) models.

For the 8600 K, log $g = 3$ model, with no stratification, we find

$$\log \left( \frac{\text{Fe}}{N_{\text{tot}}} \right) = -3.93 \text{ (Fe I)}; \quad \log \left( \frac{\text{Fe}}{N_{\text{tot}}} \right) = -3.89 \text{ (Fe II)}$$
For iron, we made separate abundance determinations for lines below the Balmer Jump. For 43 Fe I lines, we obtained log(Fe/N$_{\text{tot}}$) = $-3.86$. This is less than 0.1 dex different from the value found for lines above the Balmer Jump. Only 8 lines of Fe II were found usable below the Balmer Jump. These yielded log(Fe/N$_{\text{tot}}$) = $-4.02$. In view of the smaller number of lines, we do not consider this value significantly different from the others. All calculations of iron below the Balmer Jump were made without assuming stratification. We see no compelling evidence for stratification in iron.

The next richest spectra to Fe I and II belong to Cr I and II. We analyzed 36 Cr I lines and 47 Cr II lines. Oscillator strengths were taken from VALD (see Kupka, et al. 1999), but we omitted lines from intersystem transitions, where the multiplicity of the upper and lower terms are different. Experience has shown that ab initio and semi-empirical calculations for LS-permitted transitions are somewhat more reliable than those for intersystem lines. Reasonable agreement was found for abundances from the Cr I and II lines with wavelengths longer than $3650 \, \AA$. Only 8 lines of Cr II were found usable below the Balmer Jump. These yielded log(Cr/N$_{\text{tot}}$) = $-4.00$ (Cr I) and log(Cr/N$_{\text{tot}}$) = $-4.14$ (Cr II).

Several other elements are observed in two stages of ionization, but our results from them were less concordant. In particular, the spectra of Ti I and II may be brought into better agreement by adjusting model parameters or introducing stratification. Details are tedious, and are omitted, pending a later investigation.

### 6 Calcium

#### 6.1 The weaker Ca I and II lines

We measured equivalent widths of 23 lines of Ca I, all with wavelengths longer than the Balmer Jump. The equivalent widths spread evenly between 10 and 112 m\AA. With the 8600 K, log $g$ = 3 model, we obtained a mean log(Ca/N$_{\text{tot}}$) of $-4.65 \pm 0.08$ ($\xi_t = 3.5$). There is no significant trend of the abundance with either equivalent width or excitation potential; the latter span a narrow range of 1.88 to 2.93 eV.

The Ca I resonance line, $\lambda4227$ was not included in the above average. It is only moderately strong ($W_3 = 169$ m\AA), and its core is in not at all unusual, like those of the K-line or $\lambda\lambda4077$ or $4215$ of Sr II. We can fit the profile with log(Ca/N$_{\text{tot}}$) of $-4.65$, and lowering the $\xi_t$ to 1.5 km s$^{-1}$. We conclude $\lambda4227$ is in reasonable agreement with the weaker Ca I lines.

We found only 7 useful Ca II lines, exclusive of H and K, and the IRT. The mean value of log(Ca/N$_{\text{tot}}$) from all 7 lines was $-4.99 \pm 0.33$, with the 8600 K, log $g$ = 3 model. With these lines there were systematic trends both with excitation potential and equivalent width. Four lines, with excitation potentials averaging 7.2 eV gave log(Ca/N$_{\text{tot}}$) = $-5.5$, while three lines with average $\chi_{\text{exc}}$ of 9.0 eV gave log(Ca/N$_{\text{tot}}$) = $-4.2$. The tendency of high-excitation lines to give anomalously high abundances is a characteristic of the newly discussed anomalies of CP stars (cf. Ryabchikova, et al. 2004). However, the abundance averages from Ca I and II are within the standard errors; we have not attempted to deal with the systematic trends.

One Ca II line is below the Balmer Jump: $3461.87$ ($\chi = 7.51$). The abundance from this single line is log(Ca/N$_{\text{tot}}$) = $-4.6$. We make no judgement about stratification from this single line.

#### 6.2 The K-line

The Ca II K-line exhibits the full symptoms of stratification. An attempt to fit it with the abundance obtained from the Ca I or Ca II lines (cf. above) results in a deep, broad core, as shown in Fig. 7. The plot is quite similar to Fig. 3 of Ryabchikova, et al. (2002) for the star $\gamma$ Equi. Here, and in similar figures to follow, only token efforts were made to fit the atomic line spectrum apart from the feature of direct interest.

The stratification parameters given in the caption to Fig. 7 were determined by trial and error. They imply an...
abundance jump of 7 orders of magnitude between optical depths $\log(\tau_{5000}) = -0.7$ and $+0.3$.

### 6.3 The IRT

Lines of the IRT manifest stratification in a way that is somewhat similar to the Ca II K-line, though the total absorption is much less, and the central absorption is more significant relative to the total profile. We must keep in mind that the possible influence of surficial abundance patches which could have an influence on the profiles of the IRT.

Figure 8 shows observed and computed profiles for $\lambda 8498$ in HR 710. Damping wings are present, but the stellar profiles change rapidly in structure, from core to wing. These shapes can arise if the abundance of calcium varies with a stratification function $g(x)$ as described above.

Note the extraordinarily high value of calcium required for the deep photosphere, and the large abundance gradient implied by the parameters of the function $g(x)$.

The fit to the strongest line of the IRT, $\lambda 8542$, is shown in Fig. 9. The wings of this line are significantly stronger than those of the weaker $\lambda 8498$ line. The best fit (solid line) used different stratification parameters from those used to fit the weaker $\lambda 8498$ line (dashed line), though the abundance in the deep photosphere is roughly similar: $\log(Ca/N_{tot}) = -2.5$ vs. $-2.1$. Several lines of neutral calcium are in the region of Fig. 9. Two of the lines, at $8525.72$ and $8555.51$, were strong enough to require explicit inclusion in the calculation. The scaling factor $g(x)$ had to be applied to them, as well as $\lambda 8542$. Otherwise, with the Ca/N$_{tot}$ assumed for the deep photosphere, they would have been huge.

Figure 10 shows observed and computed profiles for $\lambda 8662$ of the IRT, as well as lower Paschen lines. The fit (solid line) used different stratification parameters from those used to fit the $\lambda 8542$ line. The dashed line is for the parameters used to fit the $\lambda 8498$ line. Both calculations use $\log g = 4$.

### 7 STRONTIUM

#### 7.1 Weaker subordinate lines

In addition to the resonance lines $\lambda \lambda 4077$ and 4215, and the stronger subordinate lines $\lambda \lambda 4416$ and 4305, we found five marginally unblended lines of Sr I and II, and determined provisional abundances from them. Results are shown in Table 1 for the 8600 K, $\log(g) = 3$ model. The abundances are up to a factor of 3 lower if the $\log(g) = 4$ model is used, but the agreement between the first and second spectra is still within 0.1 dex for the lower microturbulence. The weaker Sr I lines, which are less sensitive to microturbulence, magnetic broadening, and surface gravity give mean abundances of $-6.3$ and $-6.4$ for the lower and higher values of $\xi_t$, respectively.

We have not experimented extensively with stratification. The parameters used to get a fit for $\lambda 4077$ (see below) give unacceptably high abundances from the weak Sr I...
7.2 The resonance lines

We were unable to fit the near wings of the stronger resonance line, \( \lambda 4077 \), without moving the abundance jump deep into the photosphere. The adopted parameters given in Fig. 11 lead to \( g(x) = 0.770 \) at the deepest layer of our atmosphere, \( \log (\tau_{9000}) = 1.4 \). The strontium fraction decreases rapidly from the deepest layers, and levels to the value \( b(Sr/N_{tot}) = 2 \cdot 10^{-11} \) by \( \log (\tau_{9000}) = -5.4 \).

Fig. 12 shows the abundance stratification profiles of strontium and calcium.

The weaker of the resonance lines, \( \lambda 4215 \), is shown in Fig. 13. Thecalculation shows used the same fitting parameters as those used for the stronger line. It is unclear just how meaningful this might be, since the blending is quite significant, and extends well into the core of the profile.

7.3 The subordinate Sr \( \text{II} \) lines

Fig. 13 shows a calculation of the subordinate line, \( \lambda 4161.80 \), \( 5p^2P_{1/2}^0 - 6s^2S_{1/2} \). The line is in Sr \( \text{II} \) Multiplet 3. Only the long-wavelength part of the profile is sufficiently unblended for use in fixing the stratification parameters. The calculation shows the result of using the parameters from the best \( \lambda 4215 \)-fit, which is poor. While we were able to improve the fit in the red wing somewhat, by trial-and-error adjusting the parameters of \( g(x) \), we never achieved a satisfactory fit. When the near wing, marked ‘W’ in the figure, was fit, the calculated continuum (C) was below the observed one.

A similar situation was found for the stronger line of Multiplet 3, \( \lambda 4305 \), though for this line it was the violet wing that proved most useful (cf. Fig. 1). In all likelihood, a satisfactory fit could be found with a more elaborate stratification model than the one fixed by.

**Table 1.** Abundances from weaker subordinate Sr lines: 8600 K, log\((g) = 3.5\).

| Spectrum | \( \lambda [\text{Å}] \) | \( W_\lambda [\text{mA}] \) | \( \log (\text{Sr}/N_{tot}) \) |
|----------|----------------|----------------|-----------------|
| Sr \( \text{II} \) | 3380.69 | 97 | -6.39 \( \xi_t = 1.3 \) |
| Sr \( \text{II} \) | 3471.89 | 83 | -5.96 \( \xi_t = 1.3 \) |
| Sr \( \text{I} \) | 6408.44 | 39 | -6.21 \( \xi_t = 1.3 \) |
| Sr \( \text{I} \) | 6504.00 | 20 | -6.61 \( \xi_t = 1.3 \) |
| Sr \( \text{I} \) | 6617.27 | 18 | -5.87 \( \xi_t = 1.3 \) |
| Averages | | | -6.10 \( \xi_t = 1.3 \) |
| \( \log(\text{Sr}/N_{\odot}) \) | | | +3.02 \( \xi_t = 1.3 \) |
lated (solid black) with stratification. The parameters of the fit are: $a = 1.7$, $b = -7.0$, $d = -1$, $\log(\text{Sr}/N_{\text{tot}}) = -3.70$, the same as for the \( \lambda 4077 \) resonance line. Other atomic lines have been omitted in the calculated profile for clarity. The near wing \((W)\) and continuum \((C)\) are noted. The observed violet wing contains a weak absorption, primarily due to Fe \text{ii}. It is not strong enough to confuse the location of the near wing of the Sr \text{ii} line.

\begin{table}[h]
\centering
\begin{tabular}{cccrr}
\hline
$\lambda$[Å] & $\log(\text{El}/N_{\text{tot}})$ & $a$ & $\log b$ & $d$ & Rem. \\
\hline
Ca \text{ii} & & & & & \\
3933 & -3.10 & 6.7 & -7.00 & 0.25 & Fig. 7 \\
3933 & -4.00 & 6.7 & -6.00 & 0.50 & $\log g = 4$ \\
8498 & -2.12 & 6.7 & -5.70 & 0.25 & Fig. 8 \\
8542 & -1.52 & 1.0 & -6.70 & -1.0 & Fig. 9 \\
8602 & -1.52 & 1.0 & -6.70 & -1.0 & Fig. 10 \\
Sr \text{ii} & & & & & \\
4077 & -4.00 & 1.7 & -6.70 & -1.0 & Fig. 11 \\
4215 & -3.70 & 1.2 & -6.70 & -1.0 & See text \\
4162 & -3.70 & 1.7 & -7.00 & -1.0 & Fig. 14 \\
\hline
\end{tabular}
\caption{Summary of stratification parameters}
\end{table}

the parameters of our $g(x)$-function. We do not explore this here.

8 STRATIFICATION AND HR 710

8.1 Summary of stratification parameters

Stratification parameters are collected in Table 2. The second entries for $\lambda 3933$ give a good fit to the core and wings (to ca. ±40 Å) for a surface gravity $\log g = 4.0$. There is no corresponding figure. The entries for Sr \text{ii} 4215 are the values mentioned in the caption to Fig. 13 which give an improved fit. Note that $(\text{El}/N_{\text{tot}})$ refers to the deepest photospheric layers.

8.2 General remarks

The near wings and cores of strong lines in magnetic CP stars cannot be fit with traditional models. Non-LTE is surely relevant, but it is most unlikely to account for all of the discrepancies between theory and observation. Modern workers have returned to the notion of “anomalous atmospheres, or AA,” a term coined by Underhill (1966). Kochukhov, Bagnulo, & Barklem (2002) showed that the Balmer-line profiles could be fit with an ad hoc modification of the temperature distribution in the high photosphere. Ryabchikova and colleagues (cf. Ryabchikova 2005) have used stratified abundance models to explain both the cores and near-wings of strong lines of Ca \text{ii} and anomalous excitation and ionization, often seen in magnetic CP stars.

One might expect that the same stratification model would fit all of the Ca \text{ii} lines, H and K, and the IRT. This was not the case for HR 710. We must note that our stratification model is somewhat more constrained, by Equ. 1, than those used by Ryabchikova and coworkers.

The cores of the resonance lines of Sr \text{ii}, $\lambda \lambda 4077$ and 4215 exhibit similar peculiarities to the K-line. Vastly improved fits can be obtained using stratified models, though slightly different stratifications are required. The subordinate lines, $\lambda \lambda 4162$ and 4305 require stratification or other modifications to fit the wings. The stratification parameters required to fit $\lambda 4077$, however, make very little difference in the wings of the two subordinate lines.

Ionization and excitation anomalies are present in the HR 710 spectrum. The Ti \text{ii} and \text{ii} ionization and excitation imbalance deserves further study. The rich Fe \text{ii} and \text{ii} spectra were in satisfactory agreement without the need for stratification, provided we employed the significantly low-gravity model with $\log g = 3.0$.

Generally, the line spectrum in the red and infrared was fit better by a higher-gravity model, and the relevant illustrations shown are based on that model. We cannot rule out the relevance of abundance patches for our difficulties in fitting all of the Ca \text{ii} or the Sr \text{ii} lines with a single stratification model.

9 HR 710AA AND ITS COMPANION HR 710AB

The mass function for an SB1 system is $f(m) = m^3 \sin^3(\iota)/(m + M)^2 M_{\odot}$, where $M$ and $m$ are the masses of the primary and secondary, respectively. This $f(m)$ was well determined by Bonsack, and has the value $1.777 \cdot 10^{-3}$.

For a given $m$ and $f(m)$, the orbital inclinations decrease as $M \rightarrow m$. Kitamura (1980) noted that the inclinations of short-period Ap SB1 stars tend to be small. Such near pole-on views of SB1 systems would require a subluminous star, since the companion spectrum is unseen. Bonsack suggested a white dwarf for HR 710. This model explains the failure of the magnetic field to show an obvious reversal.

If we assume the 3-day orbital period is also the period of rotation of the primary, the observed $v \cdot \sin \iota$ depends directly on the radius of the star and the inclination of its rotational axis.

The equatorial velocity of a star is

$$v_{\text{eq}} = \frac{50.56 \cdot R}{P},$$

(2)
where \( R \) is in solar radii and \( P \) in days.

Values of \( v \cdot \sin i \) have been published by Bonsack (1981), and Preston (1971): 7.7, and 13 \text{ km s}^{-1} respectively. We have made a new determination from our UVES spectra, using the magnetic null lines \( \lambda \lambda 3850, 4006, 4065, \) and 7090, all of Fe I.

The observations were (Fourier) unfiltered, but they were interpolated to every 0.02 Å, for compatibility with Michigan software. We assumed a Gaussian 1/\( e \)-width of 0.04 Å for the instrumental profile in the blue-violet. This allows for a small degradation of the resolution; a resolving power, \( R \) of 80 000 would give ca. 0.03 Å for the 1/\( e \)-width, where we use \( \Delta \lambda = \lambda / R \) as the FWHM of the instrumental profile. For \( \lambda 7090 \), using \( R = 110 \) 000, the 1/\( e \) width of the pure instrumental profile is 0.045 Å, and we assumed 0.055 Å.

Synthetic spectra were fitted to regions containing these lines spectra, and the best fits determined by trial and error. We find that a value of 10 \text{ km s}^{-1} gives the best overall fit. This is near the mean of the Preston and Bonsack’s published values. However because of blending, and subjective judgements about a fit, the uncertainty of \( v \cdot \sin i \) is still between 1 and 2 \text{ km s}^{-1}.

We prefer \( M/M_\odot = 1.84 \) and \( R/R_\odot = 1.63 \) from tables in Drilling and Landolt (1999), for a main sequence A-star with an effective temperature of 8600 K. The values agree well with Hubrig, North, and Schöller (2007).

It is now straightforward to get the inclination from Eq. (2), if we assume a 3-day period. From the inclination, the secondary mass, \( m \), will follow from the mass function, given a primary mass, \( M \). In order to provide some indication of the overall uncertainties, we give the secondary mass in Table 3 for three assumed masses (and corresponding radii), and three values of \( v \cdot \sin i \). We assume the relation between mass and radius for main sequence stars given by Drilling and Landolt.

Our preferred values give \( m = 0.62 M_\odot \), conveniently near the most probable observational value for white dwarf masses according to Weidemann (1990). This supports the assertion of Czesla and Schmitt (2007) that the secondary of HR 710Aa “is clearly a very plausible candidate for the observed X-ray emission...”. These authors discuss a magnetically confined wind-shock model.

Bonsack (1981) also discussed a 15.88-day magnetic period. With \( R = 1.63 R_\odot \), this period leads to an equatorial velocity of 5.2 \text{ km s}^{-1}, incompatible with the adopted \( v \cdot \sin i \). To reconcile this period with the observed \( v \cdot \sin i \), we would need to assume a stellar radius nearly double that used. This is not compatible with the luminosity of the star.

**Table 3.** Secondary mass \( m \) for various primary masses \( M \) and \( v \cdot \sin i \). The preferred values are in bold face.

| \( M/M_\odot \) | \( R/R_\odot \) | \( i \) (deg) | \( v \cdot \sin i \) |
|---|---|---|---|
| 2.00 | 1.74 | 29 | 0.89 | 0.68 | 0.56 |
| 1.84 | 1.63 | 28 | 0.79 | 0.62 | 0.48 |
| 1.60 | 1.46 | 25 | 0.64 | 0.49 | 0.40 |

**10 REMARKS AND CAVEATS**

The present work on HR 710 shows that it is difficult to explain the profiles of strong Ca ii and Sr ii lines with a simple abundance-stratified model. Overall, we strongly support the stratification models, but find indications that refinements are necessary.

Different stratification parameters were required for the K-line and the IRT of Ca ii, while Sr ii lines also required adjustments of the parameters for individual lines.

It remains to be seen if theoretical work can explain the large elemental abundances of the best-fit stratification models. These require very high abundances in the deep photosphere in order to explain the observed, broad wings. In the case of calcium, traditional diffusion calculations have explained low photospheric abundances (Am stars).

Among a number of uncertainties in the present discussion, we note that our stratification model \((g(x))\) may (a) not have explored the optimum domains of parameter space, or (b) have been too simplistic. Regarding the latter point, we note the sophisticated approach of Kochukhov, et al. (2006).

Eventually, it will prove necessary to approach the modeling of the spectra of these stars with improved methods. Useful modifications of our procedures might include changing the \( T - \tau \) relation in the high photosphere. This has already been shown to give fits to the Balmer lines (Kochukhov, Bagnulo, & Barklem 2002), and can give nibs on the K-line (Cowley, Hubrig, & Kamp 2006). In the last reference, it was also pointed out that sharp K-line nibs could result if the gas pressure in the upper atmosphere were lowered.

Calculations done in LTE must be considered provisional until supported by work done with non-LTE. This is especially true for the upper atmospheres, where the total radiative flux is very nearly constant, independent of local conditions. The core of the Ca K-line is formed highest of lines considered in this work, and Cowley, Hubrig & Kamp showed that LTE calculations were only slightly modified by non-LTE.

The greatest challenge of the observations of strong-line profiles of CP stars is to the theoretician. It is straightforward to make ad hoc adjustments to a model until a profile is fit. It is quite another to justify such modifications with rigorous calculations.

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APPENDIX A: CALCULATION OF HIGHER PASCHEN LINES

Calculations near series limits need to account for overlapping lines, as well as the “dissolution” of some discrete upper levels by various perturbations from the plasma bath of the emitting and absorbing atoms. We have implemented the scheme outlined in Cowley (2000) in conjunction with a probability function $W(n)$ for the existence of discrete upper levels. The specific routine was from Barklem (2004). The basic assumption is that we can calculate the total opacity due to both bound-bound and quasi-continuous absorption by considering only a wavelength region near a given bound-bound transition. These regions extend from the halfway points between transitions to level $n-1$ and $n$, and from $n$ to $n+1$. We shall refer to these regions as the $n$-local regions. Following the reference cited, we define an $n$-dependent cross section, such that for a Paschen line

$$\int \sigma_{3n} d\Delta \lambda = \frac{\pi e^2}{mc^2} \lambda_n^2 f_{3n}. $$

The integral here is assumed to extend only over the $n$-local region. The cross sections so defined are in excellent agreement with those found in TOPbase (Cunto, et al. 1993) for photon energies below the ionization limit.

Line absorption in the $n$-local region is calculated in the usual way, but with the product $N_b f_{3n}$ diminished by the occupation probability $W(n)$. However, the line profile is truncated at each end of the $n$-local region. The ordinary continuous opacity is then augmented by an amount necessary to bring the total absorption into agreement with the above equation. In this way the total absorption merges smoothly with the known photoelectron absorption cross section at the series limit, as has been known for many years (Sugiura 1927).