CHROMOSPHERES, CORONAE AND MASS LOSS IN SOLAR AND LATE-TYPE STARS

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1. INTRODUCTION

Until observations in the EUV spectral region became available the existence of hot extended envelopes aroung latetype stars was inferred from the presence of visible and near infrared region lines which cannot be formed in radiative equilibrium within the stellar photospheres. The traditional indicators of stellar chromospheres have been the Ca II H and K lines, particularly the presence of emission components, the He I 10830 A triplet, the He II 4686 A line and emission in various hydrogen Balmer and Paschen lines. When observations from rockets and satellites become possible the strong lines of Mg II at 2800 A and the H Lyα line at 1216 A could also be included in the modelling of stellar chromospheres. The new spectra from IUE allow not only stellar chromospheres to be studied but also stellar transition regions, should hot coronae exist.

Because most of the high temperature lines in the spectral region above 1100 A are due to weak magnetic dipole transitions it is difficult to obtain direct spectroscopic evidence for material at $\sim 10^6$K from IUE observations. However, modelling based on lower temperature lines can be used to predict coronal emission measures, $\int N_e{}^2 dV$, which can be compared with measurements or upper limits available from X-ray observations.

2. Ca II AND OTHER LOW CHROMOSPHERIC LINES

Many observations of the Ca II lines have been made. Much of the interest in these lines has arisen because of the Wilson-Bappu effect (Wilson and Bappu, 1957). There remains controversy concerning the underlying physical causes of the correlation, and two main schools of thought have developed. The first, continued in recent years by Fosbury (1973), Scharmer (1976) and by Lutz and Pagel (1979), proposes that it arises through the effect of chromospheric turbulent velocites on the Doppler core of the line. The second proposes that the width is related to the damping part of the line profile, with the dependence on

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L arising from the increase in mass-column density above the temperature minimum required to give the same continuum optical depth in H^- whilst the stellar gravity decreases. Ayres, Linsky and Shine (1974), Engvold and Rygh (1978) and Ayres (1979) represent this school. Further systematic observations of Ca II line profiles have been made in recent years by Wilson (1976, 1978), Stencel (1978) and by Linsky et al. (1979).

In addition to their use in constraining models the Mg II line fluxes give directly a substantial fraction of the radiation loss above the low chromosphere, and also show a width-luminosity correlation. Following early observations from OAO-2 (Doherty 1972) and the balloon-borne spectrometer (BUSS) (Kondo et al. 1972, 1975), the Copernicus satellite has recently been used for a Mg II line survey (Weiler and Oegerle, 1979). Results from a new BUSS payload are now becoming available (Kondo et al. 1979, de Jager et al. 1979).

Earlier work on the He I 10830 A line has been extended by Zirin (1976), who finds that it is present in about 80% of all G and K stars, but only in extremely late M and S stars. IUE observations of K and M giants suggest the maximum temperature present does not exceed 2 x 10^4K.

Since it is important to determine P_o, the high chromospheric gas pressure, (where τ Ly C $\sim$ 1) various scaling laws proposed will be discussed. McClintock et al. (1975) suggested that the Mg II flux might be $\propto P_o^2/g$, (g = surface gravity). In a later paper a scaling $P_o \propto g^{\frac{1}{2}}$ is proposed, or since $P_o = m_o g$ (m_o, mass column density), $m_o \propto g^{-\frac{1}{2}}$ (Kelch et al. 1978). We note this would imply F(Mg II) = const., which is not consistent with observations, which show a strong T_{eff} dependence (Linsky and Ayres, 1978, Basri and Linsky, 1979). Moreover, when the results of Kelch (1978) for ε Eri(K2V) and 70 Oph A (KOV) are included, the fit to a $P_o \propto g^{\frac{1}{2}}$ law is poor. Using the same data, apart from α Aur (see Section 3 also), a relation F(Mg II) $\propto P_o$ fits better. Another anomaly, noted also by Kelch et al. (1979) is that the Ca II Kline flux gives a $T_{eff}^{3.7}$ dependence, compared with their T_{eff} dependence for the Mg II fluxes. Yet the work of Weiler and Oergerle (1979) leads also to F(Mg II) $\propto T_{eff}^4$. However, in spite of the uncertainties, this work showed the important result that contrary to the calculations of either Renzini et al. (1977), de Loore (1970) and Ulmschneider et al. (1977), based on acoustic heating theory, there is very little dependence of F(Mg II) on g, at given value of T_{eff}.

3. EUV EMISSION LINES

The methods needed for analyzing lines formed between 2 x 10^4 -10^5K have been developed in the context of solar work (Jordan and Wilson 1971, Burton et al. 1971), and have already been applied to α CMi by Evans et al. (1975). The starting point is the emission measure Em = $\int_R N_e^2 dh$ for each line, from which the distribution of Em with T_e can be built up (Pottasch, 1964). If each line is formed over a region R such that log T = ±0.15 dex then Em can be re-arranged to give

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$$dT/dh = P_e^2/1.4\ T_e Em \tag{1}$$

where P_e and dT/dh are taken as constant over the region of formation, but not over the whole atmosphere. P_e can be allowed to vary through the equation of hydrostatic equilibrium,

$$d \log P_e/dh = -3.1 \times 10^{-9}\ g/T_e \tag{2}$$

If the pressure P_e is known at one point then iteration between equations (1) and (2) determines the structure of the atmosphere.

The terms of the energy balance equation $\Delta F_m = \Delta F_R - \Delta F_c$ can be expressed as

$$\Delta F_R = 0.8 \int_R N_e^2\ \mathrm{Prad}\ dh \tag{3}$$

$$\Delta F_R \simeq 0.8\ Em\ \mathrm{Prad} \tag{4}$$

where ΔF_R is the radiative flux and Prad is the radiative power loss function, for example as calculated by McWirter et al. 1975. The net conduction flux ΔF_c is found from $F_c = KT^{5/2}dT/dh$ and equations (1) and (2). ΔF refers to a region corresponding to $\Delta \log T = \pm 0.15$.

If Em is expressed simply in terms of a linear fit, e.g., $Em = c\ T_e^{-2}$ (Evans et al. 1975), then for constant pressure, below $T_e = 2 \times 10^5$K,

$$F_c(T_e) = 0.7\ K\ T_e^{7/2} P_e^2/c \tag{5}$$

and $$\Delta F_c(T_e) = 2.1\ K\ T_e^{7/2} P_e^2/c \tag{6}$$

with Prad $\propto T_e^{3}$ the radiation loss becomes

$$\Delta F_R(T_e) = 8.0 \times 10^{-37}\ c\ T_e \tag{7}$$

One cannot locally set $\Delta F_R = \Delta F_c$ since the dependence of T_e is quite different. Allowing for a further term which carries energy from 2×10^5K to much lower temperatures one could justify setting the total radiation loss between 10^4 and 2×10^5K equal to the energy conducted back from the region above 2×10^5K, provided a hot corona was present. Then one would find $c \propto P_e$, at a given T_e, provided the shape of Em were the same in all atmospheres.

Although P_e can in some cases be found from chromospheric models it is useful to find P_{min}, the pressure needed to explain Em if the hottest observable line were formed in a corona.

$$P_{min} = 8.5 \times 10^{-5} (Em\ g\ T_{max})^{\frac{1}{2}}. \tag{8}$$

It is also useful to note that $\tau \propto E_m/N_e$ (Burton et al. 1971), when considering sources of line broadening.

Unless the emission measure is known above 2×10^5K (T_O) the coronal structure cannot be determined. X-ray volume emission measures help to limit possible solutions and T_C, the coronal temperature can also be measured. However a scaling law has been found (Jordan 1975, 1976) which fits quite different regions of the solar atmosphere, being based on $Em = a\, T_e^{3/2}$ above T_O. This scaling law is

$$T_c^{5/2} - T_o^{5/2} = 1.1 \times 10^8\, P_T^{\,2} T_o^{\,3/2}/Em(T_o)\, g \tag{9}$$

where P_T is the pressure at T_O. Only with an <u>additional</u> assumption of $F_R(T_O - T_C) = |F_C \text{ (at } T_O)|$ does this lead (Jordan 1979) to scaling laws similar to those found by Hearn (1975, 1977),

$$\text{i.e., } T_c^{5/2} \propto (P_T/g)^{10/9};\; E_m(T_o) \propto a \propto P_o^{8/9} g^{1/9}. \tag{10}$$

Observations of several stars on or near the main sequence are now available from IUE. The spectra of these stars are broadly similar to that of the sun, showing emission lines from both low stages of ionization such as CI, CII, OI, SiII, and also from lines formed nearer $\sim 10^5$K, e.g. CIV, SiIV, NV. The surface fluxes are comparable with or greater than those from the Sun. In order to make models and predict T_C the electron pressure is needed. Values are available for some stars through models based on CaII and MgII line profiles and fluxes. In Table 1 stars are listed for which both surface fluxes and some estimate of the pressure are available. P_O is the gas pressure at $T \sim 6500$-8000K from the models referenced. Half of P_O can be regarded as the upper limit to P_e in the transition region. Given P, g and the surface flux then P_{min} can be found from equation (8). The ratio of opacities to those in the sun also follows. The ratio of the surface fluxes to those in the sun gives immediately the ratio of the radiative losses, through equation (4). $F_C(T_O)$ can be calculated from equation (5). Finally T_C can be predicted through equation (9). For the main sequence stars it is then found that because surface fluxes are higher, but P_e is lower than in the sun, the coronal temperatures are lower than in the sun. Only if the models underestimate the pressure would these stars have coronae hotter than that of the sun.

Turning to the giant stars, in addition to αAur, αBoo and αTau, listed in Table 1, spectra are now published for μVel (G5III), εSco, αSer (K2III), (Linsky and Haisch, 1979), for αCet (M0III), (Brown <u>et al</u>. 1979) and for βGem (K0III), αTra (K2III), βAnd (M0III), and γCru (M2III) (Carpenter and Wing, 1979). Only the G giants show significant CIV emission. The later giants show low temperature species e.g. HLyα, CI, CII, SiII, SI, FeII and strong OI.

The binary star αAur (G6III + F9III) poses problems. The low pressure from the model of the primary is even less than P_{min}, and would lead to a high opacity, low temperature solution, inconsistent with the high temperature ($\sim 10^7$K) X-rays observed. The high pressure solution proposed by Baluinas <u>et al</u>. (1979), for the primary, leads to a high coronal temperature consistent with the X-ray flux.

Table 1. Observed and Derived Parameters

Star Type	log g		P dynes cm^{-2}	$\frac{E_m}{E_{m\odot}}$	$\frac{\tau}{\tau_\odot}$	T_c K	$F_c(T_c)$ erg cm^{-2} s^{-1}	Refs. and Notes
αCMi F5IV	4.2	P_o	0.055	6	17	4.1(5)	9.2(3)	Ayres et al. (1974)
		P_e	0.016			3.0(5)		
		P_{min}	0.013		38	3.0(5)	2.1(3)	Brown et al. (1979)
Sun G2V	4.4	P_o	0.16	1	1	1.5(6)	4.3(5)	
		$^*P_e$	0.077					$^*P_e$ not P_{min}
αCen A G2V	4.2	P_o	0.091	1	1.8	1.2(6)	1.5(5)	Ayres et al. (1976)
		P_{min}	4.8(-3)		16	3(5)	1.7(3)	
αCen B K1V	4.5	P_o	0.095	1	1.8	9.3(5)	1.7(5)	Ayres and Linsky (1979a)
		P_{min}	8.9(-3)		11	3(5) $^+$5(5)	5.8(3)	+X-ray T_c Nugent and Garmire (1978)
ξBoo A G 8 V K4V	4.4	^xP_e	0.13	14	8	7.9(5)	8.8(4)	^xP_e at 6300K
		P_{min}	0.023		47	2(5)	2.8(3)	Kelch et al. (1979) Hartmann et al. (1979)
εEri K2V	4.5	P_o	0.032	2.5	12	4.2(5)	7.5(3)	Linsky et al. (1978)
		P_e	0.013			2.8(5)		
		P_{min}	0.011		17	2(5)	2.5(3)	Kelch (1978)
αAur G6III +F9III	2.6	P_o	4.0(-3)	15	600	2.3(5)	19	Linsky et al. (1978)
		P_{min}	4.0(-3)		600		78	
		P_e	1.0		2.4	2(7) *1.0(7)	4.9(6)	*X-ray T_c Cash et al. (1978) Baluinas et al. (1979) Kelch et al. (1978)
αBoo K2III	1.7	P_o	1.5(-3)	$<$5(-2)	5	$^+$1.5(6)	8.2(2)	+See text
		P_e	5.6(-4)			$^+$1.1(6)		McClintock et al. (1975)
		P_{min}	1.8(4)		21	2(4)	4.8	McKinney et al. (1976) Ayres and Linsky (1975)
αTau K5III	1.8	P_o	3.5(-3)	$<$2(-3)	0.08	$^+$9.3(6)	1.1(5)	+See text
		P_e	7.6(-4)			$^+$4.8(6)		Brown et al. (1979)
		P_{min}	4.4(-5)		4	2(4)	71	Kelch et al. (1978)

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The later giants have low surface fluxes but also low pressures, and the situation regarding the coronal temperature is ambiguous. The low flux at $\sim 10^5$K may reflect a genuine absence of material, and hence $T_c \sim 2 \times 10^4$K, or a high temperature gradient, and hence rather hot coronae.

Supergiants observed with IUE include βAqr (G0Ib), αAqr (G2Ib), λVel (K5Ib) (Dupree *et al*. 1979); αPer (F5Ib) (Brown *et al*. 1979); βDra (G2II), αOri (MIab) (Linsky and Haisch 1979). In general fluxes are lower than in the sun, but early supergiants show some C IV. The K supergiants resemble the K giants. αOri is dominated by a few strong lines only and the effects of the surrounding shell have not yet been assessed.

4. MASS LOSS

The clearest forms of evidence for mass loss are an increasing outward velocity with increasing height for the emission lines, or differential expansion in circumstellar absorption lines. Although the M stars known to have mass loss do show asymmetric Mg II lines, for example in αOri and αSco, observed with Copernicus by Bernat and Lambert (1971), it is dangerous to assume the converse, that asymmetries imply mass-loss, given the variety of solar profiles observed at high spatial resolution.

The K supergiant system 32 Cyg. (K5I + BV) has been studied with IUE by Stencel *et al*. 1978. The spectra show multiple circumstellar absorption lines and lead to a mass loss of 4.7(-7) $M_\odot$/yr., consistent with earlier estimates. Most of the other recent work concerns M supergiants α^1Her (M5II), αSco A (M1.5Iab) have been observed by several groups. Reimers (1977b) and Kudritzki and Reimers (1978) find mass loss rates of 2.7(-8) $M_\odot$/yr and 7.0 (-7) $M_\odot$/yr, respectively, rather higher than the values obtained by Sanner (1976) and lower than those found by Bernat (1977). Kudritzki and Reimers now propose a mass loss relation in ($M_\odot$/yr) = 5.5(-13) L/gR (solar units). The recent BUSS payload has also been used to study circumstellar lines in αSco (Van der Hucht *et al*. 1979), from which an even higher mass loss of 7.1(-6) $M_\odot$/yr is found. Bernat's study included αOri (M2IaIb), αSco, αHer, and μCep (M2Ia), and his mass loss rates for αHer and αSco were 6.7(-7) $M_\odot$/yr and 2.2(-6) $M_\odot$/yr. Sanner's rates for these stars are about an order of magnitude smaller. Hagen (1978) has also studied a variety of M giants and supergiants, finding mass loss rates between 10(-8) and 10(-6) $M_\odot$/yr.

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