The Princeton telescope and ultraviolet spectrograph on the Copernicus satellite has been used to detect chromospheric emission lines of Lyman $\alpha$, Si $\text{III}$ and O $\text{VI}$ in the F5 IV star $\alpha$ CMi (Procyon). The interstellar absorption component of Lyman $\alpha$ leads to a H I column density of $~1.6 \times 10^{17}$ cm$^{-2}$ and a value for the neutral hydrogen density in the solar neighbourhood of $n_\text{H} = 0.015$ cm$^{-3}$. Mg II h and k emission was detected in Procyon and in the F0 Ib star $\alpha$ Car (Canopus) and the observed widths have been used to evaluate the Wilson-Bappu correlation of line width with luminosity. The analysis of the emission line intensities in Procyon leads to two models of its atmosphere in the range $10^4$ to $3 \times 10^6$ K, for the limiting cases of high and low electron pressure. In the former case a transition zone (conduction dominated regime) sets in at $10^5$ K (cf. $3 \times 10^4$ K in the Sun) indicating the existence of a corona whose temperature is deduced to lie in the range $3 \times 10^5$ to $6 \times 10^6$ K. In the latter case, the model represents a chromosphere over the temperature range observed.

I. INTRODUCTION

The emission lines associated with a chromosphere or corona in normal main sequence stars are intrinsically very weak compared with the visible photospheric continuum. However, where the stellar continuum is sufficiently depressed, e.g. in H and K of Ca II then weak chromospheric emission is often observable. The H and K emission cores have been extensively studied in stars of spectral types F to M, notably by Wilson & Bappu (1957) and by Wilson (1968). Unfortunately the line-forming processes in Ca II are dominated by non-LTE effects. Determination of chromospheric structure from H and K line profiles is extremely difficult, and in any case yields data only on the temperature range below $8 \times 10^8$ K.

The h and k lines of Mg II at 2795 and 2803 Å have been observed from balloons and space vehicles (e.g. Kondo et al. 1972) and have profiles which, for a given star, are similar to those of the Ca II H and K lines. The interpretation of their intensities is complex and in the present paper we discuss only their use in determining a Mg II ‘Wilson-Bappu’ effect. The emission lines from the higher
chromosphere, transition region and low corona (temperatures $10^4$ to $3 \times 10^5$ K) lie in the ultraviolet and are not accessible to ground-based instruments. In stars of spectral class F and later the photospheric continuum radiation is of rapidly diminishing intensity below about 2000 Å and it becomes feasible to detect the chromospheric emission lines against the much lower background. Moreover, the ultraviolet emission lines are generally optically thin and an analysis of their absolute intensities will yield directly the chromospheric structure, as in solar analyses (e.g. Jordan & Wilson 1971). In order to avoid the attenuation of the Earth's atmosphere observations from space vehicles are necessary and the results described in this paper were obtained on a guest observer basis with the Princeton Telescope-Spectrometer on the *Copernicus* satellite.

2. OBSERVATIONS

The Princeton instrumentation on *Copernicus* consists of a Cassegrain telescope and reflecting spectrograph employing a single concave diffraction grating in a Paschen-Runge mounting. The spectrum is scanned by two independent carriages designated carriage 1 and carriage 2. Each carriage contains two photomultiplier detectors, one denoted U which is sensitive to the second-order spectrum, and the other denoted V which is sensitive to the first-order spectrum. The wavelength resolution attainable with each detector is:

- $U_1 \approx 0.05$ Å, $U_2 \approx 0.2$ Å, $V_1 \approx 0.1$ Å, $V_2 \approx 0.4$ Å.

The instrument is described in detail by Rogerson et al. (1973).

Two stars of spectral type F have been observed in a guest investigator programme aimed at the detection and investigation of stellar coronae. α Car (F0 Ib, Canopus) was observed in 1973 November and α CMi (F5 IV–V, Procyon) in 1974 April. The relevant stellar parameters, taken from Hoffleit (1964) and from Allen (1973) are set out in Table I, together with the solar values for comparison.

<table>
<thead>
<tr>
<th>Star</th>
<th>Sp</th>
<th>$M_v$</th>
<th>$B-V$</th>
<th>$M$</th>
<th>$R$</th>
<th>$\log g$</th>
</tr>
</thead>
<tbody>
<tr>
<td>α CMi</td>
<td>F5 IV</td>
<td>$+2.65$</td>
<td>$+0.41$</td>
<td>$1.77 M_\odot$</td>
<td>$2.0 R_\odot$</td>
<td>$4.09$</td>
</tr>
<tr>
<td>α Car</td>
<td>F0 Ib</td>
<td>$-4.7$</td>
<td>$+0.16$</td>
<td>$12 M_\odot$</td>
<td>$60 R_\odot$</td>
<td>$1.95$</td>
</tr>
<tr>
<td>Sun</td>
<td>G2 V</td>
<td>$+4.83$</td>
<td>$+0.65$</td>
<td>$1 M_\odot$</td>
<td>$1 R_\odot$</td>
<td>$4.44$</td>
</tr>
</tbody>
</table>

Both stars were studied with observing programmes designed to detect particular emission lines which are strong in the upper chromosphere and transition zone of the Sun, taking due account of the spectral region where the *Copernicus* sensitivity is high, i.e. in the range 1000–1300 Å. Because of the low expected intensity of the stellar emission lines, each spectral region was scanned many times and, after inspection to remove obviously unreliable data, the scans were averaged to produce the mean spectrum.

In Canopus the only unambiguous emission was in the Mg II h and k lines, $\lambda 2795$ and 2803 Å, shown in Fig. 1. The emission is very weak and appears as a distortion of the photospheric absorption, the probable photospheric profile being shown as a broken line. In Procyon the Mg II emission is stronger and it has been observed from a rocket-borne experiment of Kondo et al. (1972).
3. ABSOLUTE FLUX CALIBRATION

The data obtained from the Copernicus instruments are in the form of a photoelectron count rate as a function of wavelength and it is necessary to convert this to an absolute flux measurement. A suitable method is to compare Copernicus observations of a particular star with observations of the same star by other absolutely calibrated instruments. The bright hot star ξ Puppis is a suitable transfer standard since its ultraviolet spectrum has been observed by many different instruments, in particular by those used on the rocket flights by Stecher (1970) and Burton et al. (1973), and by the S2/68 sky survey telescope described by Boksenberg et al. (1973). Of these three instruments, S2/68 and Stecher’s provided good (∼30 per cent) photoelectric photometry but at a low wavelength resolution (10–35 Å); the results of Burton et al. were of much higher wavelength resolution (0.3 Å) but the photometry was not as good (∼50 per cent) as a result of the photographic method of recording the spectrum. The intercomparison was performed in spectral regions where the high resolution rocket data showed there to be little structure in the spectrum over at least 10 Å so that errors introduced by the different wavelength resolutions would be minimized. The high resolution rocket spectrum was also used to identify stellar features with essentially zero residual flux, which made it possible to estimate the stray light level in the Copernicus U2 scans which were known to be seriously contaminated by scattered light.

The resulting absolute calibration is estimated to be accurate to better than
Fig. 2. Copernicus observations of chromospheric emission lines in Procyon: (a) Lyman \( \alpha \) 1216 Å; (b) Si \( \text{II} \) 1206.5 Å; (c) O \( \text{VI} \) 1032 Å.

© Royal Astronomical Society • Provided by the NASA Astrophysics Data System
The calibration constants \( \eta(\lambda) \) (expressed as \( \eta(\lambda) = (\text{photons cm}^{-2} \text{2}^{-1} \text{Å}^{-1})/(\text{U2 counts in 14 s}) \)) and the derived absolute fluxes received at the Earth are set out in Table II.

**Table II**

<table>
<thead>
<tr>
<th>Ion</th>
<th>( \lambda ) (Å)</th>
<th>( \eta(\lambda) )</th>
<th>Observed flux (erg cm(^{-2}) s(^{-1}))</th>
<th>Surface flux (erg cm(^{-2}) s(^{-1}))</th>
</tr>
</thead>
<tbody>
<tr>
<td>O vi</td>
<td>1032</td>
<td>0.020</td>
<td>( 1.5 \times 10^{-13} )</td>
<td>( 2 \times 10^3 )</td>
</tr>
<tr>
<td>Si iii</td>
<td>1206</td>
<td>0.050</td>
<td>( 1.4 \times 10^{-12} )</td>
<td>( 8.4 \times 10^3 )</td>
</tr>
<tr>
<td>H i</td>
<td>1216</td>
<td>0.045</td>
<td>( 1.9 \times 10^{-11} )</td>
<td>( 6 \times 10^3 )</td>
</tr>
<tr>
<td>O v</td>
<td>1218</td>
<td>0.045</td>
<td>(&lt; 2 \times 10^{-13} )</td>
<td>( 2 \times 10^5 )</td>
</tr>
<tr>
<td>C ii</td>
<td>1335</td>
<td>0.243</td>
<td>(&lt; 1.3 \times 10^{-12} )</td>
<td>( 5 \times 10 )</td>
</tr>
</tbody>
</table>

* \( H i \) \((N_H = 3 \times 10^{17} \text{cm}^{-2})\).
† \( H i \) \((N_H = 1.5 \times 10^{17} \text{cm}^{-2})\).

4. **INTERPRETATION AND ANALYSIS**

(a) *Mg ii h and k emission*

The H and K lines of Ca ii are accessible to ground-based observations and have been extensively studied as indicators of stellar chromospheric activity. The most remarkable result is the strong correlation of emission line width with stellar luminosity, the Wilson–Bappu effect. Kondo et al. (1972) have observed profiles of chromospheric Mg ii emission in three stars, including Procyon, and Moos et al. (1974) have determined the Mg ii profiles in Arcturus. The present *Copernicus* observations of Canopus and the observations of solar Mg ii, for instance by Purcell, Boggess & Tousey (1961), now provide a total of six stars with known Mg ii emission line widths.

In Fig. 3 the width of Ca ii emission is plotted against stellar absolute magnitude \( M_v \), and the straight line through the Ca ii observations represents the Wilson–Bappu effect. Also plotted are the Mg ii observations and, although these points are fewer in number and show a greater scatter than those of Ca ii, it is possible to say that the observations are quite consistent with a ‘Wilson–Bappu’ type relation for Mg ii, but with a line width consistently 2.6 times larger than for Ca ii.

The observation and interpretation of chromospheric emission line widths have been discussed by Fosbury (1973) who derives a line width for Mg ii which is 1.8 times larger than that of Ca ii. The difference between this ratio and the one derived above is due to the fact that Fosbury's measurements refer to the FWHM, which is essentially equivalent to the widths measured photographically by Wilson & Bappu, while the present measurements refer to the width close to the base of the emission profile.

(b) *Interstellar H i absorption*

The observed Lyman \( \alpha \) profile (Fig. 2) clearly contains a component due to absorption by neutral hydrogen in the line of sight to Procyon. Although a certain amount of self reversal is to be expected in the intrinsic line profile, it would be
difficult to conceive of a stellar atmospheric structure which would reduce the residual flux in the line core essentially to zero. The interstellar absorption in Lyman $\alpha$ is sufficiently strong that the interstellar line is formed mainly in the radiation damped wings of the transition. In these circumstances the relation between the equivalent width $W$ of the absorption and the column density of interstellar neutral hydrogen $N_H$ is independent of the velocity dispersion of the absorbing atoms, and is given by

$$N_H = 1.865 \times 10^{18} W^2 \text{ atom cm}^{-2}, \text{ with } W \text{ in Å.}$$

In the case of the Procyon Lyman $\alpha$ profile it is not possible to calculate $W$ since the intrinsic profile of the chromospheric emission line is not known. It is, however, possible to calculate the interstellar H I column density by a rather different procedure.

In the damping wings of Lyman $\alpha$ the optical depth $\tau(\lambda)$ due to a column density $N_H$ atom cm$^{-2}$ is given by $\tau(\lambda) = 9.25 \times 10^{-20} (\lambda - \lambda_0)^{-2} N_H$, where $\lambda_0$ is the wavelength of the line centre, and $\lambda$ and $\lambda_0$ are in Å. The intrinsic stellar emission at wavelength $\lambda$ will then be attenuated by a factor $e^{-\tau(\lambda)}$ in propagating through the interstellar hydrogen. If a value of $N_H$ is assumed then the intrinsic stellar emission profile may be reconstructed by multiplying the observed flux by $e^{\tau(\lambda)}$. Since the single observed point at the line centre has zero flux it is necessary to interpolate this point from its immediate neighbours. For consistency, the interpolated point, when multiplied by $e^{-\tau(\lambda)}$ must return to essentially zero flux.

Two of these reconstructed profiles are shown in Fig. 4, corresponding to $N_H = 1.6 \times 10^{17} \text{ cm}^{-2}$ and $3.0 \times 10^{17} \text{ cm}^{-2}$. These are close to the lower and upper
limits of \(N_H\) that will produce a reasonable profile for the intrinsic stellar emission, observations of solar Lyman \(\alpha\) being used as a guide to the expected profile. The distance to Procyon is 3.5 pc (Allen 1973) and the mean density of neutral hydrogen in the line of sight is thus \(0.015 \text{ cm}^{-3} < N_H < 0.03 \text{ cm}^{-3}\). This value of \(n_H\) is low compared with \(n_H \approx 0.3 \text{ cm}^{-3}\) deduced by Macchetto & Panagia (1973) and by Jenkins & Savage (1974) both for much more distant stars, at distances of about 100 pc. A similar low value of \(n_H = 0.015 \text{ cm}^{-3}\) has been calculated by Moos et al. (1974) for the 11 pc line of sight to \(\alpha\) Boo (Arcturus) and it would thus appear that the value of the hydrogen density in the immediate solar neighbourhood, i.e. within about 10 pc of the Sun, is only \(\approx 0.015 \text{ cm}^{-3}\).

This low value of \(n_H\) implies that the opacity of the nearby interstellar medium at wavelengths below the \(\text{H} \text{I}\) Lyman limit (\(\lambda < 912 \text{ Å}\)) is less than was previously

![Graph of reconstructed intrinsic stellar Lyman \(\alpha\) profiles](image)

**Fig. 4.** Reconstructed intrinsic stellar Lyman \(\alpha\) profiles: (a) \(N_H = 1.6 \times 10^{17} \text{ cm}^{-2}\); (b) \(N_H = 3.0 \times 10^{17} \text{ cm}^{-2}\). The ordinate is an arbitrary linear flux scale.
estimated (Aller 1959). According to the calculations of Crudace et al. (1974) a mean density of $n_H = 0.015 \text{ cm}^{-3}$ and a path length of 10 pc would result in approximately 10 per cent transmission at the Lyman edge, and the transmission would exceed 50 per cent at all wavelengths shorter than the He I ionization limit ($\lambda = 504 \, \text{Å}$).

The detection of coronal soft X-ray emission in nearby stars is thus limited solely by the intrinsic stellar emission and not by interstellar absorption and these stars would make ideal candidates for observations in the as yet unexplored 100–912 Å region.

(c) Absolute fluxes at the stellar surface

The analysis of the chromospheric structure of Procyon requires first that the observed emission line fluxes be transformed into fluxes at the stellar surface. This requires a knowledge of the radius and distance of the star or, alternatively, of its angular diameter. Fortunately the angular diameter of Procyon has been measured by an intensity autocorrelation interferometer (Hanbury Brown et al. 1967) and the angular diameter of the equivalent uniform disk is given as $\theta_{\text{UD}} = 5.31 \times 10^{-3}$ arcsec. The effect of limb darkening is to change this by only a few per cent and this has been ignored. The flux at the stellar surface $F_s$ is then related to the flux at the Earth $F_e$ simply by

$$\frac{F_s}{F_e} = \left( \frac{2}{\theta_{\text{UD}}} \right)^2 = 6 \times 10^{15}.$$  

Table II also sets out the calculated emission fluxes at the surface of Procyon with the solar values for comparison.

(d) Analysis of emission measures

The absolute fluxes observed can be used to derive the emission measure $\int_R N_{e_2}^2 \, dh$ for each line, using methods developed for previous analyses of solar data (e.g. Pottasch 1964). Assuming that, as in the Sun, the observed emission is formed by collisional-excitation from the ground state, the flux at the star can be expressed as

$$F = 7.8 \times 10^{-16} f \bar{g} g_m(T) \frac{N(E)}{N(H)} \int_R N_{e_2}^2 \, dh \, \text{erg cm}^{-2} \text{s}^{-1}$$  

(1)

where $f$ is the oscillator strength, $\bar{g}$ is the Gaunt factor, $N(E)/N(H)$ is the element abundance, $g_m(T)$ is the maximum value of the function

$$g(T) = T_{e_2}^{1/2} \frac{N(\text{ion})}{N(E)} 10^{-5040 W_{12}/T_e}$$  

(2)

$N_{e_2}, T_{e_2}$ are the electron density and temperature respectively and $R$ is the region of the atmosphere over which the line is formed.

Table III gives the atomic data used in the analysis, and for the lines of C II, Si III and O VI these are the same as used by Burton et al. (1971), in their analysis of solar intensities.

Fig. 5 shows the values of the emission measure derived from the Procyon data, and for comparison, the solar values, and the mean curve from Burton et al. (1971). The Lyman $\alpha$ points have been calculated assuming all photons created by collisional excitation from the ground state eventually escape from the atmosphere in...
the observed line. If processes other than spontaneous radiative decay from the excited level are important the derived emission measures represent lower limits. For Si III it has been assumed that all the ions are in the ground state. If the metastable triplet levels had a Boltzmann population, only one-third of the ions would be in the ground state. In view of the known transition probability for the decay of the metastable level and the probable low density of the atmosphere of Procyon (see below) the former approximation is likely to be the more appropriate. The O V limit is too high to be useful, and was found by assuming collisional excitation to $2s2p^3P_0,1,2$ is balanced by radiative decay of $2s2p^3P_1$. The collision strength of Eissner (see Gabriel & Jordan 1972) was used to find the effective $gf$ product.

If a smooth curve is drawn through the observed points then the distribution of $\int N_e^2 \, dh$ with temperature can be used to find a range of models of the atmosphere which would satisfy the observations. The method of analysis derived by Jordan

![Graph](image-url)

**Fig. 5.** The emission measures $\int N_e^2 \, dh$ for Procyon as a function of electron temperature $T_e$. The corresponding solar values and the solar mean curve are also shown.
(1965) and as applied by Burton et al. (1971) has been used. In this method the observed distribution of $\int_{R} N_{e}^{2} \, dh$ gives

$$\int_{R} N_{e}^{2} \, dh = \int_{\log T_{1}}^{\log T_{2}} N_{e}^{2} \frac{dh}{dT_{e}} \cdot d \log T_{e}$$

(3)

where $\log T_{1}$ and $\log T_{2}$ are $\log T_{m} \pm 0.15$.

Assuming, initially, that $dh/d\log T$ and $P_{e} = N_{e}T_{e}$ are constant over $\Delta \log T$,

$$\int_{R} N_{e}^{2} \, dh = \frac{[P_{e}^{2}]_{\Delta T}}{T_{e}^{2}} \cdot \left[ \frac{dh}{d \log T} \right]_{\Delta T} \times 0.32.$$  

(4)

Making the assumptions of hydrostatic equilibrium and a fully ionized gas gives

$$\frac{d \log P_{e}}{d \log T_{e}} = -0.39 \times 10^{-4} \frac{T_{e}}{P_{e}} \frac{d \log T_{e}}{d \log T}$$

(5)

where the surface gravity of Procyon has been taken as $1.2 \times 10^{4}$ cm s$^{-2}$. This is derived from a mass of $1.77 M_{\odot}$ (Allen 1973) and a radius of $1.98 R_{\odot}$ (Hanbury-Brown et al. 1967).

Then starting from a given pressure at a given temperature, the pressure gradient and temperature gradient can be found through the iteration of equations (4) and (5). In the solar atmosphere the boundary condition for the pressure in the corona is known, enabling a narrow range of solutions to be found, typically giving $P_{e} = 5.6 \times 10^{14}$ cm$^{-3}$ K (Jordan 1966). Unfortunately no such boundary condition is known for stellar atmospheres and the range of permissible solutions is much larger.

Although the pressure at $T_{e} \sim 3 \times 10^{5}$ K is not known, the total gas pressure in the photosphere and low chromosphere has been calculated in a model of Procyon made by Ayres, Linsky & Shine (1974). At the top of the chromosphere ($T_{e} = 8000$ K) Ayres et al. find a total gas pressure of $3.4 \times 10^{14}$ cm$^{-3}$ K. This figure can therefore be used to give an upper limit to the electron pressure at $T_{e} = 3 \times 10^{5}$ K (where the medium is fully ionized) of $P_{e} = 1.7 \times 10^{14}$ cm$^{-3}$ K. A lower limit to the electron pressure at $3 \times 10^{5}$ K can be found by assuming that the O VI emission is formed in an isothermal medium of this temperature. The scale height in such a medium would be, from equation (6), $H = 3.3 \times 10^{4}$ km. Combining this with the absolute value of the emission measure for O VI leads to $[N_{e}^{2}]^{1/2} = 6.8 \times 10^{7}$ cm$^{-3}$ and $P_{e} = 2.0 \times 10^{13}$ cm$^{-3}$ K. Fig. 6 shows the variation of pressure with temperature derived using these limiting values, together with the model of Ayres et al. (1974).

The models derived for the upper and lower limits in electron pressure are tabulated in Tables 4(a) and (c) respectively. In the latter case of low electron pressure the structure derived from the $\int N_{e}^{2} \, dh$ distribution is significantly different from that in the Sun in that the temperature gradient decreases steadily between $T_{e} \sim 1.5 \times 10^{4}$ K and $T_{e} \sim 3 \times 10^{5}$ K, whereas in the Sun this gradient passes through a maximum in that region.

In the case of the higher pressure model, it will be seen that a conductively dominated region and, hence, a corona does exist but further knowledge of the shape of the emission measure distribution is required in order to determine its structure. Suppose that above $T_{e} \sim 3 \times 10^{5}$ K the emission measure increases according to $aT^{3/2}$, as is often assumed for the Sun, where $a$ is determined by the emission measure at $3 \times 10^{5}$ K. Then, using also the equation of hydrostatic
in the chromosphere of Procyon, also showing the model of the low chromosphere by Ayres et al. (1974).

equilibrium it can be shown (Jordan 1974) that the coronal temperature is related to $P_0$, the electron pressure at $T_0$ (in this case $3 \times 10^5$ K) by

$$T_e^{5/2} = \left( \frac{P_0^2}{a} \right) \frac{1}{D^2}$$

where $D = 3/H$, $T_e = \text{const}$ and $H$ is the scale height.

Hence using $P_0^2$ and $a$ at $3 \times 10^5$ K gives

$$T_e = 1.3 \times 10^6 \text{ K}.$$  

If the emission measure continued to fall at $T_e > 3 \times 10^5$ K before increasing, again according to the $aT^{3/2}$ power law, then even higher temperatures would be allowed. It is therefore important to determine in more detail the shape of the emission measure distribution between $T_e \approx 5.6 \times 10^4$ K and $3 \times 10^5$ K to find whether the O vi line is formed where the emission measure is increasing or decreasing.

(e) The energy balance in Procyon's atmosphere

In order to facilitate calculations simple analytical forms will be used for the emission measure and radiative losses. From Fig. 5 it can be seen that the emission measure can be represented by

$$\int R N_e^2 \, dh = cT^{-2}$$  

where $c = \text{const} = 1.0 \times 10^{36}$. Hence from equations (3), (4) and (7),

$$P_0^2 \left( \frac{dh}{d \log T} \right) = 3.13 \times 10^{36} = \text{const}.$$
Using also equation (5) gives

$$P_e^2 = P_0^2 + 2 \cdot 44 \times 10^{32} (T_e^{-1} - T_0^{-1})$$  \hspace{1cm} (9)

where $P_0$ is $P_e$ at $T_0 = 3 \cdot 10^5$ K.

Further, from McWhirter, Thonemann & Wilson (1974), between $T_e \sim 10^4$ K and $3 \times 10^5$ K the radiative loss function is given by

$$P_{\text{rad}} = \frac{d}{T^{4/5}} \text{erg cm}^{-3} \text{s}^{-1}$$  \hspace{1cm} (10)

where $d = 3 \cdot 2 \times 10^{-26}$.

The energy transported by conduction is given by

$$F_c = 1 \cdot 0 \times 10^{-6} T^{5/2} \frac{dT}{dh} \text{erg cm}^{-2} \text{s}^{-1}$$  \hspace{1cm} (11)

and this can be calculated immediately for the models given in Table IV.

Now assume that over each temperature region $\Delta h$ corresponding to $\Delta \log T_e = 0.1$, the energy dissipated is balanced by the sum of the radiation losses,

**Table IV**

*Models derived from the observations*

(a) Model 1—using mean curve

<table>
<thead>
<tr>
<th>$\log T_e$</th>
<th>$\log N_e$</th>
<th>$\log P_e$</th>
<th>$\log (dh/d \log T_e)$</th>
</tr>
</thead>
<tbody>
<tr>
<td>4.2</td>
<td>10.14</td>
<td>14.34</td>
<td>7.70</td>
</tr>
<tr>
<td>4.3</td>
<td>10.03</td>
<td>14.33</td>
<td>7.77</td>
</tr>
<tr>
<td>4.4</td>
<td>9.92</td>
<td>14.32</td>
<td>7.89</td>
</tr>
<tr>
<td>4.5</td>
<td>9.80</td>
<td>14.30</td>
<td>7.96</td>
</tr>
<tr>
<td>4.6</td>
<td>9.69</td>
<td>14.29</td>
<td>8.03</td>
</tr>
<tr>
<td>4.7</td>
<td>9.58</td>
<td>14.28</td>
<td>8.15</td>
</tr>
<tr>
<td>4.8</td>
<td>9.47</td>
<td>14.27</td>
<td>8.23</td>
</tr>
<tr>
<td>4.9</td>
<td>9.36</td>
<td>14.26</td>
<td>8.29</td>
</tr>
<tr>
<td>5.0</td>
<td>9.26</td>
<td>14.26</td>
<td>8.31</td>
</tr>
<tr>
<td>5.1</td>
<td>9.15</td>
<td>14.25</td>
<td>8.33</td>
</tr>
<tr>
<td>5.2</td>
<td>9.04</td>
<td>14.24</td>
<td>8.37</td>
</tr>
<tr>
<td>5.3</td>
<td>8.94</td>
<td>14.24</td>
<td>8.35</td>
</tr>
<tr>
<td>5.4</td>
<td>8.83</td>
<td>14.23</td>
<td>8.35</td>
</tr>
<tr>
<td>(5.5)</td>
<td>8.73</td>
<td>14.23</td>
<td>8.56</td>
</tr>
</tbody>
</table>

(b) Model 1—using analytical forms

<table>
<thead>
<tr>
<th>$\log T_e$</th>
<th>$\log N_e$</th>
<th>$\log P_e$</th>
<th>$\log (dh/d \log T_e)$</th>
<th>$F_c$</th>
<th>$\Delta F_c$</th>
<th>$- \Delta F_r$</th>
<th>$\Delta F_m$</th>
</tr>
</thead>
<tbody>
<tr>
<td>4.2</td>
<td>10.12</td>
<td>14.32</td>
<td>7.88</td>
<td>15.2</td>
<td>6.51 \times 10^4</td>
<td>6.51 \times 10^4</td>
<td></td>
</tr>
<tr>
<td>4.3</td>
<td>10.01</td>
<td>14.31</td>
<td>7.90</td>
<td>32.5</td>
<td>3.76 \times 10^4</td>
<td>2.85 \times 10^4</td>
<td></td>
</tr>
<tr>
<td>4.4</td>
<td>9.90</td>
<td>14.30</td>
<td>7.92</td>
<td>69.5</td>
<td>3.29 \times 10^4</td>
<td>2.16 \times 10^4</td>
<td></td>
</tr>
<tr>
<td>4.5</td>
<td>9.78</td>
<td>14.28</td>
<td>7.96</td>
<td>1.42 \times 10^8</td>
<td>72.5</td>
<td>3.75 \times 10^4</td>
<td>2.13 \times 10^4</td>
</tr>
<tr>
<td>4.6</td>
<td>9.68</td>
<td>14.28</td>
<td>7.96</td>
<td>3.18 \times 10^8</td>
<td>1.76 \times 10^3</td>
<td>1.64 \times 10^4</td>
<td>1.57 \times 10^4</td>
</tr>
<tr>
<td>4.7</td>
<td>9.56</td>
<td>14.26</td>
<td>8.00</td>
<td>6.47 \times 10^8</td>
<td>3.29 \times 10^3</td>
<td>1.24 \times 10^4</td>
<td>1.07 \times 10^4</td>
</tr>
<tr>
<td>4.8</td>
<td>9.45</td>
<td>14.25</td>
<td>8.02</td>
<td>1.39 \times 10^8</td>
<td>7.43 \times 10^3</td>
<td>1.64 \times 10^4</td>
<td>1.57 \times 10^4</td>
</tr>
<tr>
<td>4.9</td>
<td>9.35</td>
<td>14.25</td>
<td>8.02</td>
<td>3.11 \times 10^8</td>
<td>1.72 \times 10^3</td>
<td>1.24 \times 10^4</td>
<td>1.07 \times 10^4</td>
</tr>
<tr>
<td>5.0</td>
<td>9.24</td>
<td>14.24</td>
<td>8.04</td>
<td>6.63 \times 10^8</td>
<td>3.52 \times 10^3</td>
<td>9.43 \times 10^3</td>
<td>5.91 \times 10^3</td>
</tr>
<tr>
<td>5.1</td>
<td>9.14</td>
<td>14.24</td>
<td>8.04</td>
<td>1.49 \times 10^4</td>
<td>8.27 \times 10^3</td>
<td>7.15 \times 10^3</td>
<td>-1.12 \times 10^3</td>
</tr>
<tr>
<td>5.2</td>
<td>9.04</td>
<td>14.24</td>
<td>8.04</td>
<td>3.34 \times 10^4</td>
<td>1.85 \times 10^4</td>
<td>5.43 \times 10^3</td>
<td>-1.31 \times 10^3</td>
</tr>
<tr>
<td>5.3</td>
<td>8.93</td>
<td>14.23</td>
<td>8.06</td>
<td>7.11 \times 10^4</td>
<td>1.77 \times 10^4</td>
<td>4.12 \times 10^3</td>
<td>-3.36 \times 10^3</td>
</tr>
<tr>
<td>5.4</td>
<td>8.83</td>
<td>14.23</td>
<td>8.06</td>
<td>1.59 \times 10^6</td>
<td>8.79 \times 10^3</td>
<td>3.12 \times 10^3</td>
<td>-8.48 \times 10^3</td>
</tr>
<tr>
<td>(5.5)</td>
<td>8.73</td>
<td>14.23</td>
<td>8.06</td>
<td>3.57 \times 10^8</td>
<td>1.98 \times 10^5</td>
<td>2.36 \times 10^3</td>
<td>-1.96 \times 10^5</td>
</tr>
</tbody>
</table>
Observations of emission lines in F stars

Table IV—continued

(c) Model 2—using mean curve

\[
\begin{array}{cccccc}
\log T_e & \log N_e & \log P_e & \log (dh/d \log T_e) \\
4.2 & 9.91 & 14.11 & 8.14 \\
4.3 & 9.78 & 14.08 & 8.26 \\
4.4 & 9.65 & 14.05 & 8.42 \\
4.5 & 9.51 & 14.01 & 8.54 \\
4.6 & 9.37 & 13.97 & 8.67 \\
4.7 & 9.23 & 13.93 & 8.85 \\
4.8 & 9.08 & 13.88 & 9.01 \\
4.9 & 8.93 & 13.83 & 9.16 \\
5.0 & 8.76 & 13.76 & 9.29 \\
5.1 & 8.60 & 13.70 & 9.44 \\
5.2 & 8.42 & 13.62 & 9.61 \\
5.3 & 8.23 & 13.53 & 9.75 \\
5.4 & 8.03 & 13.43 & 9.95 \\
(5.5 & 7.81 & 13.31 & 10.24) \\
\end{array}
\]

(d) Model 2—using analytical forms

<table>
<thead>
<tr>
<th>\log T_e</th>
<th>\log N_e</th>
<th>\log P_e</th>
<th>\log (dh/d \log T_e)</th>
<th>\Delta F_e</th>
<th>\Delta F_r</th>
<th>\Delta F_m</th>
</tr>
</thead>
<tbody>
<tr>
<td>4.2</td>
<td>9.90</td>
<td>14.10</td>
<td>8.32</td>
<td>5.52</td>
<td>-</td>
<td></td>
</tr>
<tr>
<td>4.3</td>
<td>9.75</td>
<td>14.05</td>
<td>8.42</td>
<td>9.89</td>
<td>4.4</td>
<td>6.51 x 10^4</td>
</tr>
<tr>
<td>4.4</td>
<td>9.60</td>
<td>14.00</td>
<td>8.52</td>
<td>17.5</td>
<td>7.6</td>
<td>4.93 x 10^4</td>
</tr>
<tr>
<td>4.5</td>
<td>9.44</td>
<td>13.94</td>
<td>8.64</td>
<td>29.7</td>
<td>12.2</td>
<td>3.75 x 10^4</td>
</tr>
<tr>
<td>4.6</td>
<td>9.29</td>
<td>13.89</td>
<td>8.74</td>
<td>52.7</td>
<td>23.0</td>
<td>2.85 x 10^4</td>
</tr>
<tr>
<td>4.7</td>
<td>9.14</td>
<td>13.84</td>
<td>8.84</td>
<td>93.6</td>
<td>40.9</td>
<td>2.16 x 10^4</td>
</tr>
<tr>
<td>4.8</td>
<td>8.98</td>
<td>13.78</td>
<td>8.96</td>
<td>1.59 x 10^3</td>
<td>65.4</td>
<td>1.64 x 10^4</td>
</tr>
<tr>
<td>4.9</td>
<td>8.82</td>
<td>13.72</td>
<td>9.08</td>
<td>2.69 x 10^3</td>
<td>1.10 x 10^3</td>
<td>1.24 x 10^3</td>
</tr>
<tr>
<td>5.0</td>
<td>8.67</td>
<td>13.67</td>
<td>9.18</td>
<td>4.81 x 10^3</td>
<td>2.12 x 10^3</td>
<td>9.43 x 10^3</td>
</tr>
<tr>
<td>5.1</td>
<td>8.50</td>
<td>13.60</td>
<td>9.32</td>
<td>7.80 x 10^3</td>
<td>2.99 x 10^3</td>
<td>7.15 x 10^3</td>
</tr>
<tr>
<td>5.2</td>
<td>8.34</td>
<td>13.54</td>
<td>9.44</td>
<td>1.32 x 10^3</td>
<td>5.40 x 10^3</td>
<td>5.43 x 10^3</td>
</tr>
<tr>
<td>5.3</td>
<td>8.16</td>
<td>13.46</td>
<td>9.60</td>
<td>2.05 x 10^3</td>
<td>7.30 x 10^3</td>
<td>4.12 x 10^3</td>
</tr>
<tr>
<td>5.4</td>
<td>7.99</td>
<td>13.39</td>
<td>9.74</td>
<td>3.34 x 10^3</td>
<td>1.29 x 10^3</td>
<td>3.12 x 10^3</td>
</tr>
<tr>
<td>(5.5</td>
<td>7.81</td>
<td>13.31</td>
<td>9.90</td>
<td>5.15 x 10^3</td>
<td>1.81 x 10^3</td>
<td>2.36 x 10^3</td>
</tr>
</tbody>
</table>

the thermal conduction and stellar wind losses, i.e.

\[
\Delta F_m + \Delta F_r + \Delta F_c + \Delta F_w = 0
\]  

(12)

where

\[
\Delta F = \int_{\Delta h} \frac{dF}{dh} dh,
\]

\( F \) is an energy flux and the subscripts \( m, r, c \) and \( w \) refer to mechanical, radiative, conductive and stellar wind terms respectively. Initially, the latter will be taken as negligible. The radiative losses are given by

\[
\Delta F_r = - \int_{\Delta h} 8 \cdot 80 N_e^2 P_{\text{rad}} dh
\]  

(13)

using \( N_H = 8 \cdot 80 N_e \).

Substitutions from equations (8) and (10) lead to

\[
\Delta F_r = - 8 \cdot 80 \times 10^9 T^{-6/5}
\]  

(14)
Hence, the energy balance equation (12) may be written as

$$\Delta F_m = 8.0 \times 10^9 T^{-6/5} - \Delta F_c.$$  (15)

The parameters of the models derived from the simplified analytical expressions (8) and (9) for the two limiting cases of high and low electron pressure are listed in Tables 4(b) and (d) respectively. Also given are the conductive ($\Delta F_c$) and radiative ($\Delta F_r$) dissipations together with the corresponding value of dissipation of mechanical energy ($\Delta F_m$) required to form the energy balance in the absence of other energy losses such as a stellar wind.

It can be seen that for the low-density-limit model thermal conduction never dominates, and the energy balance is between mechanical energy dissipation (probably shock heating) and radiative losses, i.e. the atmosphere can be regarded as a chromosphere which extends to much higher temperatures than in the case of the Sun.

For the high pressure model, the dissipation of conductive flux becomes dominant at $T_e = 2 \times 10^5$ K indicating the presence of a transition zone and, hence, of a true corona at greater heights. Indeed the negative dissipation in mechanical flux indicates that there is too great a dissipation in conductive flux to be accommodated by radiation. A similar energy problem can exist in the case of the Sun and may be due to an inhomogeneous atmosphere structure (Kopp & Kuperus 1968). Another possibility is that the stellar wind term makes up the deficiency and, although this is not important for the quiet Sun, it is of interest to calculate the magnitude of a stellar wind for Procyon which would be needed to establish an energy balance in this model (see Section 4(f)).

It would appear from the present results that the chromosphere of Procyon extends to a significantly higher temperature than in the Sun since the onset of a conductively dominated region occurs at a higher temperature than in the Sun. However, the presence of a true corona needs further investigation and depends on the actual value of electron pressure in the upper atmosphere on the star. An upper limit cannot be derived directly from the observations and was therefore estimated from the following overall energy consideration. This assumes that the form of the mechanical energy flux is an acoustic wave. Then the maximum energy flux that the medium can carry is that which corresponds to a Mach one shock wave, i.e.

$$F_m \leq \frac{1}{2} \rho C_s^3$$

where $\rho$ is the local density and $C_s$ is the local sound velocity. If we now take the high pressure model (Table 4(a)), an upper limit to the energy flux available for heating the corona is given by applying the above inequality to the base of the transition zone, i.e. at $T_e = 10^5$ K. This gives $F_m \leq 1.4 \times 10^5$ erg cm$^{-2}$ s$^{-1}$.

For acoustic heating, an approximate analytical solution for a coronal model has been derived by McWhirter et al. (1974) which can be expressed in terms of the heating energy flux as follows

$$F_m = 3.2 \times 10^{-13} P_e T_e^{1/2}.$$  

Use of the above limit in $F_m$ gives the upper limit in coronal temperatures to be

$$T_e \leq \left( \frac{F_m}{3.2 \times 10^{-13} P_e} \right)^2 = 6 \times 10^6$$

This argument depends on a number of assumptions but some support is afforded
by the corresponding value in the Sun derived for the model of Jordan (1965).
This is $T_c \leq 2 \times 10^6$ K, a value which is not far above the generally accepted value
($1.6 \times 10^6$ K) for the solar corona. It is likely that the upper limit in energy flux is
rather close to the true value, explaining the close agreement for the Sun. For
Procyon, a further upper limit is imposed by the adopted limit in electron pressure.

(f) The effect of a stellar wind on the energy balance

If a stellar wind exists then the transport of material through a temperature
gradient leads to an energy loss rate in erg cm$^{-3}$ s$^{-1}$

$$W = \frac{d}{dh} \left\{ v \frac{2}{3} N k T + \frac{1}{2} N m v^2 \right\}$$

where $v$ is the local stellar wind velocity, $N$ is the particle density, $m$ is the mean
particle mass.

In order to conserve matter $N v = \text{const}$ and we obtain

$$W = N v \left\{ \frac{3}{2} k T \frac{d}{dh} (\ln T) - m v^2 \frac{d}{dh} (\ln N) \right\}.$$ 

In an atmosphere where the pressure $NT$ is approximately constant

$$\frac{d}{dh} (\ln T) = - \frac{d}{dh} (\ln N)$$

so that as long as the stellar wind velocity is much less than the sound speed

$$W \simeq \frac{3}{2} N v k T \frac{d}{dh} (\ln T).$$

In order to compare this energy loss with Table 4(b) we need $\int W \, dh$, the integral
being over $\Delta T = 0.1$ dex. Thus

$$\Delta F_W = - \int W \, dh = - \int \frac{3}{2} N v k T \frac{d\ln T}{dh} \, dh$$

$$= -0.345 \, N v k T.$$

For the stellar wind to balance the energy discrepancy of Table 4(b) at $\log T = 5.3$,
we must have $\Delta F_W \sim -3 \times 10^4$ cm$^{-2}$ s$^{-1}$.

Thus

$$N v \simeq 2.5 \times 10^{15} \text{ cm s}^{-1}.$$ 

$N$ is about equal to twice the electron density, thus

$$v \simeq 3.7 \times 10^6 \text{ cm s}^{-1} \equiv 37 \text{ km s}^{-1}.$$ 

This implies a stellar mass loss rate $\dot{M} = 1.5 \times 10^{15} \text{ g s}^{-1}$ ($\equiv 2.2 \times 10^{-11} \, M_\odot \, \text{yr}^{-1}$)
compared with the mass loss due to the solar wind $\dot{M} \simeq 8.6 \times 10^{10} \text{ g s}^{-1}$.
Thus the stellar wind of Procyon would need to be $\sim 2 \times 10^4$ as intense as the solar wind.

The theory of the generation of stellar winds is not adequately understood and
it is impossible to state whether or not this relatively intense stellar wind is reasonable
for a star of the spectral class of Procyon. However, both the lower surface
gravity and increased luminosity of Procyon relative to the Sun would suggest a
more intense stellar wind. A stellar wind velocity of 37 km s\(^{-1}\) should be detectable spectroscopically by its Doppler effect on line profiles and this is discussed in Section (g) below.

(g) Emission line widths

The observations of chromospheric emission line widths were made with the U2 detector which has an exit slit corresponding to \(0.2 \, \text{Å}\), and for which the detector step increments are \(0.2 \, \text{Å}\). The signal to noise ratio in the observations of Si \(\text{III}\) and O \(\text{VI}\) emission lines is low but the fact that more than one point in the spectral scans of these lines was significantly above background does enable some estimates of line widths to be made.

In the case of O \(\text{VI}\) two adjacent points are significantly above background but this does not lead to a useful lower limit on the line width since the line may have been observed near to the edge of the \(0.2 \, \text{Å}\) slit. The upper limit to the line width corresponds to a full width at half maximum (FWHM) of about \(0.4 \, \text{Å}\).

In the Si \(\text{III}\) observations there are three adjacent points which are significantly higher than the background which would imply a FWHM of the Si \(\text{III}\) line of about \(0.4-0.6 \, \text{Å}\). Poisson statistics cannot be used to determine the confidence level of this result since there is an electronic prescaling device in the detector system. However, an analysis of the fluctuations in the observed counts suggests that the probability of the Si \(\text{III}\) line width being at least \(0.4 \, \text{Å}\) FWHM is greater than 0.75. It is clearly dangerous to draw any important conclusions from this result, but since future observations should determine the true width of the Si \(\text{III}\) emission, it is useful to examine the possible consequences of a line width of \(0.4 \, \text{Å}\), which is much greater than the width of \(0.05 \, \text{Å}\) caused by thermal motions at the Si \(\text{III}\) line forming temperature of \(55000 \, \text{K}\).

The effect of the stellar rotational velocity \(v \sin i = 6 \, \text{km s}^{-1}\) is negligible but the observed line width could be produced by radiative transfer if the optical depth in the line were much greater than unity. However, using the models given in Table IV the calculated optical depths are too small to produce the observed widths.

Boland et al. (1973) have considered the emission line widths produced by a propagating acoustic wave in the solar chromosphere and their analysis shows that an acoustic wave of Mach number \(\alpha = 1\) would produce a FWHM of \(0.27 \, \text{Å}\) in \(\lambda 1206.5\) if the line were optically thin. Since the observed line width is greater than this, the required Mach number of an acoustic wave is greater than unity. Such a wave would be very strongly damped and would not propagate in the manner required to heat the chromosphere and corona. This problem is not alleviated by invoking Alfvén or magneto acoustic waves since the observed line width is determined by the harmonic motion of the ions and in any propagating wave this velocity \(v\) cannot exceed the local sound speed. The energy transported by the wave is, however, dependent on the mode of oscillation and for an acoustic wave of Mach number \(\alpha = 1\) this amounts to \(\frac{1}{2}Nm\bar{v}^2 = \frac{1}{2}Nmcs^2 = 6.0 \times 10^4 \, \text{erg cm}^{-2} \, \text{s}^{-1}\).

A stellar wind is another possible line broadening mechanism due to the projection effects of a radial streaming motion when viewing the complete stellar disk. However, in order to conserve mass, \(\rho v = \text{constant}\), and since the density \(\rho\) must be lower in the O \(\text{VI}\) line forming region than in the Si \(\text{III}\) line forming region, then the velocity must be higher at O \(\text{VI}\) than at Si \(\text{III}\). Thus any stellar wind that is
adequate to produce the observed line width in Si III will produce a larger Doppler broadening in O VI and this is contrary to the observations. It must be concluded that a simple radially flowing stellar wind cannot explain the observations.

The intrinsic nature of the difficulty in explaining the width of $\lambda\,1206.5$ is that the inferred Doppler velocity ($50\,\text{km}\,\text{s}^{-1}$ half width at half maximum intensity) is supersonic. Any such motion will be rapidly damped unless the motion is coherent with a scale length much greater than the damping length. This of necessity rules out all conceivable wave or turbulent motions, but does allow large scale mass motions, possibly convective in origin, or a non-radial stellar wind.

In view of the relatively low confidence in the Si III line width these conclusions must be regarded as tentative.

5. CONCLUSIONS

The profile of the Lyman $\alpha$ emission line of Procyon has been used to derive a column density of neutral hydrogen in the line of sight of $N_H \approx 1.6 - 3 \times 10^{17}$ atom $\text{cm}^{-2}$. This column density is sufficiently low as to attenuate radiation from Procyon below the Lyman limit by at most 75 per cent. Procyon is thus potentially observable at all wavelengths in the far-ultraviolet and soft X-ray regions.

Besides Lyman $\alpha$ the resonance lines of Si III and O VI have been observed in emission in Procyon, and this is the first time that these ions have been observed in stars of spectral class F. The calculated absolute intensities have been used to obtain the emission measures $\int N_e^2\,dh$ for Si III, O VI and also Lyman $\alpha$.

Methods previously used to analyse the solar chromospheric spectrum allow the density and temperature profile of the atmosphere of Procyon to be determined over the temperature range $10^4$ to $3 \times 10^6$ K, given that the boundary value of electron pressure is known. This latter parameter cannot be determined directly from the present observations but limiting values can be established to give $1.7 \times 10^{14} \geq N_eT_e \geq 2.0 \times 10^{13}$ cm$^{-3}$ K. The models derived for these two limiting cases have been examined for energy balance. The high pressure model shows the existence of a true transition zone, i.e. a region where conductive flux is dominant, and hence of a corona at greater heights. However, the transition zone is thicker and its onset occurs at a higher temperature ($10^5$ K) than in the case of the Sun ($3 \times 10^4$ K). The low pressure model shows no such transition zone and is best regarded as an extended chromosphere. It appears that the chromosphere of Procyon extends to greater heights and higher temperatures than in the Sun but the presence of a corona will need further investigations. Limits to the temperature of any corona are deduced to give $3 \times 10^6 \, K < T_c < 6 \times 10^6$ K.

Further observations are desirable in order to refine the present models of the chromosphere and transition region and in particular observations below 912 Å are needed to determine the coronal structure.

ACKNOWLEDGMENTS

We would like to thank Dr Lyman Spitzer for the generous provision of guest observer time on the Princeton Telescope-Spectrometer on Copernicus and also the staff of Princeton University Observatory, especially Dr D. G. York and Dr T. P. Snow for their assistance in planning and reducing the Copernicus observations.
REFERENCES