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Physics of the Solar Atmosphere [and Discussion]

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Physics of the solar atmosphere

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A summary is given on recent results on the physics of the quiet solar atmosphere, and active regions. This includes: solar rotation, velocity fields and waves, magnetic field concentration, the transition region, coronal magnetic field structure, and prominences.

1. Solar rotation

The main problems are:

(*a*) has the Sun a rigidly rotating core?
(*b*) how to explain the observed differential rotation?
(*c*) is there a secular decrease of the solar angular momentum?
(*d*) why do the magnetic regions rotate faster than the non-magnetic?

(*a*) *A rigidly rotating core*

The existence of a rigidly rotating core has been proposed on many occasions in the past, and based on different arguments. The most direct observations which, for the first time, gave observational evidence for the existence of such a core are the following:

The magnetic photospheric regions rotate faster than the non-magnetic ones, by 0.1 km/s at the equator (for a review see Stenflo 1974*b*) and therefore show less differential rotation than the non-magnetic regions. This proves that the magnetic regions are anchored to another level of the solar body with a different rotational velocity, and shows in any case that there is a depth gradient of solar rotation.

Coronal holes, which often extend from the equator to a pole possess an overall angular velocity that is more or less independent of latitude (Timothy, Krieger & Vaiana 1975).

Combined with the first-mentioned observation this shows that the coronal holes are anchored to a *rigidly* rotating part of the Sun. This indicates the existence of an inner core and confirms earlier, more indirect inferences, based on calculations of the internal circulation and viscosity. The core should rotate slightly faster than the envelope. A very fast spinning core proposed by Dicke (see review, 1970) is not confirmed by these observations.

(*b*) *Differential rotation*

Figure 1, after Stenflo (1974*a*) summarizes our knowledge of the differential rotation of the Sun. Interesting is the difference between the magnetic and non-magnetic parts.

The dashed, solid and dotted lines correspond to increasing values of the flux density. The angular velocity of the photospheric plasma is smaller than that of the photospheric field by 0.1 km/s (at the equator). Stenflo interpreted this by the assumption that the magnetic elements are very small, and that an eastward streaming plasma flows *around* the small magnetic elements. Larger elements such as sunspots offer a larger dynamic pressure than the non-magnetic plasma flow and are thus partly dragged along. This picture still needs to be explained quantitatively.

There are two groups of theories to explain the differential rotation.

(i) In their *simplest* form the Boussinesq theories assume a fast rotating ellipsoidal rigid core (Jeans 1926; De Jager 1959, p. 343). Convection transports angular momentum into a radial direction, and since for any freely moving convective element $\omega^2 r$ should remain constant, ω should vary as r^{-2} which would explain the differential rotation.

(ii) It is known for a long time already that by the interaction of rotation with convection meridional currents must develop in the convection zone, and authors like Ward (1965), Roxburg (1970) and Durney (1974*a*, *b*) showed how the action of the Coriolis forces on the meridional circulation would produce differential rotation.

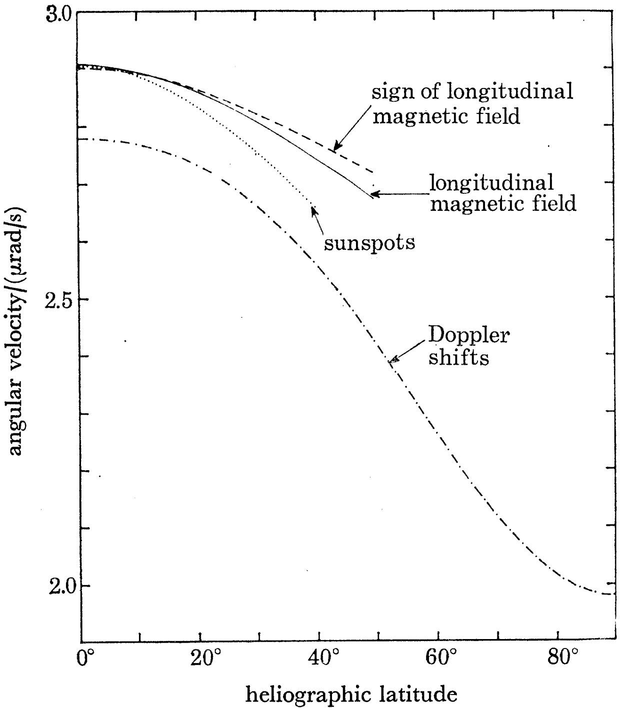

FIGURE 1. Solar rotation (after Stenflo 1974*a*).

(*c*) *Decay*

Solar angular rotation will decay for several reasons. According to Van den Heuvel & Conti (1971), the *e*-folding time of solar rotation is approximately 2.2×10^9 years (a). For the rotational energy this time is half this value, being 1.1×10^9 a. Since the solar rotational energy is 2×10^{35} J, the dissipation rate is 2×10^{18} J s^{-1}.

Whether this loss of solar angular momentum can be explained by the solar wind alone is not yet clear: on the basis of the observed parameters of the solar wind Brandt & Heise (1970) found an *e*-folding time of $3-4 \times 10^9$ a.

2. VELOCITY FIELDS AND WAVES

Convection

We summarize briefly: The outer solar convection layer extends to a monochromatic optical depth at 5000 Å, $\tau_5 \approx 0.8$, but convective elements arriving at that level may have excess heat and momentum, so that the upward motions continue till $\tau_5 \approx 0.08$. The up- and downward velocities are 1–2 km/s (velocity *differences* of 2–4 km/s, close to the sound speed). In the convective regions the more or less systematic granular motions develop a field of turbulent motions on smaller scales. There are indications (Rutten, Hoyng & de Jager 1974) that this velocity field has a Kolmogoroff spectrum, but this observation needs further confirmation.

The convection region is the source of a mechanical energy flux; its value at the basis of the chromosphere ($\tau_5 \approx 10^{-3}$) is approximately 0.2 J cm^{-2} s^{-1}.

Wave modes

The motions in the convective region can be described by a complex of waves of different modes depending on the restoring forces: gas pressure (compression waves), gravitation (gravity waves), and magnetic tension and/or pressure (Alfvén waves of magnetic–acoustic waves). If gas pressure and gravitation are both acting, the modes of the acoustic–gravity waves can best be described by a diagnostic dispersion diagram giving the boundaries of the regions where pure acoustic or pure gravity waves can propagate vertically (figure 2 from the review by Stein & Leibacher 1974). Since the dispersion diagram gives relations between the periods and wavelengths there is – in principle – a possibility of observationally identifying wave modes. However there are drawbacks.

First of all the observations refer to a region very near the source regions so that from that point of view the k–ω diagram is an unsuitable tool for the analysis of solar photospheric motion modes; furthermore the observational possibilities are very limited due to lack of spatial resolution. Secondly, the observed inhomogeneity in the photosphere influences strongly the existence, the propagation and the nonlinear interaction of the different modes, all of which are still hardly understood.

As things stand now it seems that compression (acoustic) waves propagating out of the convective region are mainly responsible for the heating of the low chromosphere (Ulmschneider 1975).

3. CONCENTRATION OF MAGNETIC FIELDS IN THE QUIET PHOTOSPHERE AND CHROMOSPHERE

Magnetic flux concentrations are known to exist at the photospheric and low chromospheric level. These concentrated flux tubes necessarily fan out at higher levels since the gas pressure decreases with height faster than the magnetic pressure. Hence, from a certain height upwards the magnetic pressure exceeds the gas pressure. For a B-value of 100 G (10^{-2} T) the critical level occurs at a height of 150 km above the limb ($\tau_5 \approx 2 \times 10^{-4}$); for stronger fields like those occurring in spots, and perhaps, in the filigree elements the level would still be lower. The magnetic structures at the coronal and upper chromospheric level must be more diffuse than in the photosphere, since large differences in the magnetic pressure cannot be compensated by the gas pressure.

This picture is actually confirmed by the observation of the chromospheric fine structures made in emission lines emitted at different heights of the chromosphere–coronal transition region and beyond. The disappearance of the small scales of the chromospheric fine structure at higher levels shows the fanning out of the magnetic field pattern. At these heights the magnetic field determines the structure completely; inhomogeneities can be expected on scales comparable with the height above the photosphere, due to the equilibrium between the magnetic pressure and tension. Some concentration of the fields remains in existence up into the corona.

Characteristic field concentrations are assumed to be: the filigree; the field at the network boundaries (the coarse and fine mottles of the network); and the X-ray bright points.

We shall not discuss here the flare pores (supra-thermal plasma nodules) and the sunspots.

Magnetic field structure of the network

Already on visible inspection the network is seen to consist of individual elements, the coarse mottles, which fall apart into still smaller structures, the fine mottles. The latter have sizes of approximately 1″ or smaller, but the true diameter cannot be determined for instrumental and atmospheric reasons. The network boundary coincides with a concentration of the magnetic field and it seems likely that the field is restricted to the mottles.

In this connexion we mention the filigree (Dunn & Zirker 1973) which is related to a concentration of magnetic fields in photospheric and low chromospheric elements. Although there is not much support for the initial hypothesis of Livingston & Harvey (1969) that the flux seems quantized, there are indications that the peak values of magnetic field strengths in these small elements are large. Zwaan (1967) already suggested these to be of the order of 0.1 T and this was confirmed by Stenflo (1973). This has now been observed as the filigree structure (Frazier, this volume, p. 295; Rust, this volume, p. 427, 1976). The sizes of these elements should typically be in the range of 100–300 km. These magnetic field concentrations have also been discussed theoretically: Peckover & Weiss (1972) treated the concentration of magnetic field in the presence of convection and found that locally the magnetic field strength can be much larger than the equilibrium value; recently Parker (1974*a*, *b*) discussed the convective pumping to explain the magnetic flux concentrations.

X-ray bright points

These structures (Vaiana 1976, this volume, p. 365; figure 3) belong to the most enigmatic results of rocket X-ray photography, and were confirmed and studied in more detail by the A.t.m. observers. They are very small bright features, randomly and uniformly distributed over the disk and seem to correspond with small bipolar field regions.

In Hα and Ca K they correspond to bright patches. However, most bright spots in Ca K are not coronal bright points. Areas range from 10^7 to 2×10^8 km^2. Observations with A.t.m. (Golub *et al.* 1974) showed them to have average life times of 8 h, statistically distributed. It is estimated that about 1500 X-ray bright points emerge per day. They seem to be identical with small 'newly emerging flux regions' studied by Harvey & Martin (1973). Golub *et al.* (1974) estimate from the equipartition argument, using T and N determinations from spectral data that $B \approx 10^{-3}$ T. Then the flux should be approximately 10^{11} Wb and with 1500 points originating daily, this would lead to a total emerging flux of 10^{14} Wb/day. We should add immediately, contrary to the implicit assumption of Golub *et al.*, that this should not mean a loss of magnetic flux of that order: if the points are bipolar and remain so the flux should for the greater part submerge again at the end of the life-time of the bright point.

The emergence of flux loops

The problem how the flux tubes originate and appear at the solar surface is one of the most important in solar physics, and in stellar astrophysics as well. With regard to the *appearance* of flux tubes at the solar surface Parker's (1955) ideas still stand: the tubes are lifted by buoyancy forces. The lateral motion in the supergranules is thought to produce field concentration at the boundaries of the cells, although Glackin (1974) finds no relation between the emergent flux

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regions and the network. On the other hand it is commonly assumed that the strong fields formed at the boundary of the network *are* formed by the outward streaming motions in the cells. Then the main problem is how the photospheric flux tubes, with fields of 10^{-1}–10^{-2} T can still exist in higher layers.

Independently, several authors have come to the only possible solution: the tubes must be bent or twisted and thus contain a large amount of energy. They loose their energy by Ohmic dissipation which leads to heating and emission of X-radiation, and appear as the bright points (Parker 1974*a*; Stenflo 1974*a*; Piddington 1974; Glackin 1974). In details as well as in the quantitative elaboration these discussions differ, but that aspect does not seem important for present discussion.

4. THE TRANSITION REGION BETWEEN THE CHROMOSPHERE AND THE CORONA

The transition region is characterized by an extremely large temperature gradient essentially caused by the inability of the upper layers to radiate the dissipated mechanical energy in a low temperature configuration. It has been realized for a long time that heat conduction must play a major role because of the high temperature and the large temperature gradients. Early e.u.v. data lead to conduction dominated models of the transition region with fluxes in between $5\times10^{-2}\,\mathrm{J\,cm^{-2}\,s^{-1}} < F_\mathrm{c} < 1.2\times10^{-1}\,\mathrm{J\,cm^{-2}\,s^{-1}}$. In all these models it has been assumed that the transition region is in hydrostatic equilibrium and homogeneous, so that every part of the solar disk contributed equally to the intensity of an emission line. Moreover it has been assumed that the pressure was almost constant throughout the transition region. The basic parameters that determined these empirical models were T_cor, p and F_cond. In a fully ionized plasma the coefficient of heat conduction is $K = \kappa T^{\frac{5}{2}}$, where $\kappa = 5\times10^{-14}\,\mathrm{J\,cm^{-1}\,s^{-1}\,K^{-\frac{7}{2}}}$. In the case of a hydrostatic atmosphere the structure of a plane transition layer is determined by the energy balance equation

$$q_\mathrm{h} + q_\mathrm{cond} = q_\mathrm{r}, \tag{1}$$

where q_h is the mechanical heating, q_r the radiative losses and $q_\mathrm{cond} = \mathrm{d}(\kappa T^{\frac{5}{2}}\,\mathrm{d}T/\mathrm{d}h)/\mathrm{d}h$ the conductive heating. Introducing the quantity $\eta = \frac{2}{7}\kappa T^{\frac{7}{2}}$ the energy equation becomes

$$\mathrm{d}^2\eta/\mathrm{d}h^2 = q_\mathrm{r} - q_\mathrm{h}. \tag{2}$$

An analysis of this equation is given by Lamers & Kuperus (1974). It is evident that a steep temperature rise levelling off in an almost constant temperature corona requires that η has an inflexion point. At that point $q_\mathrm{r} = q_\mathrm{h}$, while for higher levels the heating dominates and for lower levels the radiation losses dominate. It is customary to use the relation given by Cox & Tucker (1969) for the radiative losses. Above $T = 10^5$ K the radiation losses decrease rapidly. Therefore the η-inflexion point should be located at a level h_i in the atmosphere, where the temperature is higher than 10^5 K. It is not too bad an approximation to neglect q_r beyond this level. Then the temperature structure can be found simply by integrating equation (2) remembering that $q_\mathrm{h} = -\mathrm{d}F/\mathrm{d}h$, where F is the mechanical energy flux.

A rough estimate of the maximal coronal temperature can be found under the assumption that F decays exponentially with an *e*-folding length $L \ll d$, where d is the distance between the level of maximum coronal temperature and the inflexion point. It then follows that

$$\eta_\mathrm{m} \approx F_\mathrm{i} L \tag{3}$$

and the maximum coronal temperature is only determined by the mechanical energy flux at the inflexion point and the dissipative damping length L. If we insert for $L = 5 \times 10^8$ cm and for $F_1 = 10^{-1}\,\mathrm{J\,cm^{-2}\,s^{-1}}$ we find $T_{\max} = 1.4 \times 10^6$ K in good agreement with the observed coronal temperature (Lamers & Kuperus 1974).

However, plane hydrostatic models seem to be too much of an idealization to interpret recent observations made with OSO-VII and Skylab instruments. In the first place there is strong evidence for large amplitude motions comparable with the thermal velocities (Boland, Engstrom, Jones & Wilson 1973). In that case it certainly is not correct anymore to assume that the pressure is constant throughout the whole transition region. It has been shown by Flower & Pineau des Forets (1974) that, taking the dynamic pressure ρv^2 into account, a constant density model is more appropriate than a constant pressure model and that the conductive flux is certainly not constant in the transition region.

A different reason why the atmosphere cannot be purely in hydrostatic equilibrium has been given by Kuperus & Athay (1967). The chromosphere is already radiating at maximum efficiency and how should the atmosphere dispose of the energy conducted backwards? If the atmosphere could expand for instance through spicule motions the conducted energy could be balanced. Bessey & Kuperus (1970) demonstrate that impulsive heating could result in spicule type motions in the solar chromosphere, while Browne & Bessey (1973) showed how an equilibrium transition region develops giving rise to a systematic outflow consistent with the solar wind observations.

Homogeneous models were useful in as far as they gave us the feeling to work with the energy balance in a medium with a very large temperature gradient, but the transition layer is very inhomogeneous. Kopp & Kuperus (1968) drew the attention to the fact that in a magnetic field heat is conducted along the field lines thus setting up a temperature distribution along the field lines. If, moreover, most of the coronal magnetic field lines are anchored in photospheric magnetic concentrations it is expected that, due to the channelling of the heat into these magnetic knots (located at the supergranulation boundary), the structure of the transition layer above the supergranulation boundary deviates markedly from that above the centre of the cells, which has been confirmed by observations. Kopp (1972) argued that since the e.u.v. emission originates from about 20 % of the solar surface the conductive flux inferred from the e.u.v. line intensities is much smaller than previously determined values. Therefore the heat deposited in the upper chromosphere could be radiated and even a considerable mechanical heating is required to balance the radiative losses.

An inhomogenous model of the transition region has been constructed by Gabriel (1976, this volume).

The A.t.m. results show that the e.u.v. emission above the supergranulation boundary spreads out when measured in lines that are formed higher in the transition layer. It seems that the e.u.v. emission strongly reflects the presence and the intensity of magnetic fields. This leads to the important conclusion that the mechanical energy flux that reaches the lower parts of the transition region is strongly dependent on the magnetic field strength. It is not clear whether the field only affects the transport of mechanical flux through the temperature minimum or whether it is also crucial in the generation of the mechanical flux itself. The generation of Alfvén and/or magneto acoustic waves in the photosphere by the turbulent convection, the nonlinear interaction between sound waves and these m.h.d. waves, and the propagation of these waves in an inhomogeneous structure should be thoroughly investigated. Since in general, in a low β plasma the m.h.d. waves are less damped than in a high β plasma, it is clear that the magnetic concentrations can channel mechanical energy flux in the form of m.h.d. waves into the transition region with greater ease than in neighbouring regions with low magnetic field strength.

This is consistent with the observation that in active regions the e.u.v. emission is enhanced over the whole area of the active region, while in coronal holes, which are assumed to be the regions of very weak coronal fields, the e.u.v. emission is strongly reduced or almost absent (Noyes & Withbroe 1972). Withbroe & Gurman (1973) give as representative values of temperature for a hole, a normal corona and an active region, 10^6 K, 1.6×10^6 K and 2.5×10^6 K respectively (see also Goldberg 1974).

5. Coronal magnetic fields

The corona is a very inhomogeneous hot plasma consisting of a number of structures like streamers, loops, polar plumes, coronal holes and coronal condensations above the active regions.

The only way to understand the structure of the solar corona is to assume that relatively strong magnetic fields are present. If we assume that 5 % of the photospheric surface is covered with magnetic concentrations of the order of 0.1 T and if we moreover assume that within a layer of the thickness of the chromosphere and the transition region these fields spread out so that the magnetic flux is homogeneously distributed in the inner corona the magnetic field strength should then be of the order of 5×10^{-4} T. At one solar radius distance the field strength $B \approx 1 \times 10^{-4}$ T.

Actually the coronal field strengths are somewhat smaller since a considerable amount of magnetic flux is present in closed loops and most of them are far below one solar radius height. Let us now compare the magnetic pressure with the gas pressure. The gas pressure in the corona varies from 10^{-3} Pa at $R_\odot$ to 10^{-4} Pa at $2R_\odot$. Hence $\beta = p_{\text{gas}}/p_{\text{mag}} < 1$, while nearly everywhere in the photosphere $\beta > 1$. This change of β from a large value in the photosphere towards a small value in the corona is the prime reason for the fact that the degree of inhomogeneity in gas pressure increases when going from the photosphere into the corona. For small values of β large fluctuations in the gas pressure can be maintained. Moreover since the conduction of heat occurs essentially along the magnetic field, neighbouring regions with quite different temperatures may exist.

In what respect could the magnetic field in the corona be determined from the observed structure? Many attempts have been made to compute the coronal field structure using the observed photospheric magnetic field distribution and then compare the calculated field lines with structures visible in the white light corona (Altschuler & Newkirk, 1969).

These calculations show that a potential field distribution is in good agreement with the loop type structures and the streamers seen on the photographs. A similar result was obtained by Rust & Roy (1971) who found that above an active region a system of potential magnetic field lines could be found that coincided perfectly with a system of loop prominences. Now the results of these potential field calculations should only be considered as a qualitative agreement. For instance a force free field could as well fit the observed structures even when a considerable amount of energy is stored in the field. Recently Levine & Altschuler (1974) superposed a system of electric currents on a large scale calculated potential field and mapped the new configuration. It was found that large currents are required to get significant topological deviations from the potential magnetic field configuration. Their conclusion is that any agreement between coronal structures and calculated potential field configurations should not be interpreted as proof that currents are insignificant.

Indeed electric currents can be significant in certain coronal regions. Since the corona is a low β plasma a non-potential magnetic flux tube can only be maintained if the magnetic field is twisted or put differently if a current flows along the magnetic field. It is still an open question why certain flux tubes are visible in X-ray emission where others are not (cf. Jordan 1976, this volume, p. 391).

Magnetic flux tubes may either emerge from the subphotospheric layers where they have already received an appreciable amount of twisting, or they may originate in the corona by the local evolution of coronal magnetic fields. In the first case we do not expect that the fields are *a priori* stable against magnetohydrodynamic instabilities. Because the field strains have formed in the subphotospheric high β plasma it is likely that the field strongly expands and rises when entering the low β coronal plasma. In the second case existing coronal fields must be twisted presumably by photospheric motions of the field line footpoints. A possible way to generate a current with a component parallel to the magnetic field is to rotate one of the footpoints of the magnetic field with respect to the other.

The axial current induced by these motions, grows as long as the field lines are being wound up. If this occurs in a low β plasma, the azimuthal magnetic field set up by the axial component of the currents constricts the plasma and counteracts the gradient in the pressure of the axial field component until there is force balance. The result is an almost force free coronal filament with a density enhancement. It is of importance for the evolution of coronal magnetic fields to investigate whether these magnetic flux tubes are stable against magnetohydrodynamic and thermal disturbances.

A dense flux tube looses more energy by radiation than the less dense ambient medium. For optically thin coronal plasma the radiation losses are given by

$$q_{\mathrm{r}} = \Phi(T)\rho^2/\mathrm{J\,cm^{-2}\,s^{-1}}, \tag{4}$$

where the function $\Phi(T)$ is given by Cox & Tucker (1969). If we now assume that the radiation losses are completely balanced by the heating of the corona, q_{h}, for every stable flux tube as well as for the ambient corona it follows that

$$\frac{\Phi(T_{\mathrm{i}})}{\Phi(T_{\mathrm{e}})} = \frac{\rho_{\mathrm{e}}^2 q_{\mathrm{hi}}}{\rho_{\mathrm{i}}^2 q_{\mathrm{he}}}, \tag{5}$$

where subscripts (i, e) denote the interior of a flux tube and the exterior, respectively. It is unlikely that the heat input is just proportional to the density squared and thus flux tubes with different densities must also have different temperatures.

Inspection of the curve $\Phi(T)$ given by Cox & Tucker around a temperature of about 10^6 K shows that a density difference of a factor two, and thus a ratio of $\Phi(T_{\mathrm{i}})/\Phi(T_{\mathrm{e}}) = 0.25$ can be fulfilled if $T_{\mathrm{i}} \approx 2T_{\mathrm{e}}$. So considerable temperature differences are expected in the coronal flux tubes.

If on the other hand $q_{\mathrm{hi}} \gg q_{\mathrm{he}}$ while the density ratio is about the same as in the above example the temperature inside the flux tube has to decrease so that $\Phi(T_{\mathrm{i}})$ increases and the excess heating can be compensated. It is then quite natural that where excessive heating takes place $\Phi(T)$ will be maximal, when the density is fixed otherwise. This means that $T \approx 10^5$ K as can be seen from the radiative loss function. This could explain the highly twisted coronal filaments observed in He II 304 Å from Skylab.

It should be mentioned here that heating and cooling of a low β plasma does not effect the density very much. We thus expect the corona to be filled with flux tubes of varying density and quite different temperatures, the density being determined by the amount of twisting of the magnetic field for example.

Although most of the corona has a low value of β there are regions where β is large or at least of the order of unity. An example of such a region is the coronal magnetically neutral sheet where the field strength is very small. Any heating or cooling results in an expansion or a compression of the medium. The magnetic tension may be too small to resist the expansion. If this occurs the initially closed field lines may burst open to form a coronal streamer. As soon as an open magnetic structure is present the coronal plasma can escape freely into interplanetary space thus creating an additional energy loss. This means that the coronal temperature in these regions cannot be as high as in the closed field regions. Again this seems to be confirmed by recent observations of the coronal holes where we observe a depression of X-ray emission above magnetically very quiet regions and where an enhancement of the solar wind flux is anticipated (Krieger *et al.* 1973; Noci 1973).

If in the high β regions a suppression in the heating or a slight increase in the cooling occurs this does not only result in a subsequent cooling as it would in the low β regions as well but also in an associated compression resulting in a density increase. The increased density combined with the decreased temperature causes the medium to cool further and it is thus expected that a thermal instability occurs leading to the formation of prominences. We will concentrate on this aspect of coronal magnetic fields in the subsequent section.

6. PROMINENCES

Solar prominences are cool filamentary structures embedded in the corona showing a large variety in morphological structure and life times approximately in between 10^3 and 10^6 s. The longlasting quiescent prominences which occur between regions of opposite magnetic polarities differ remarkably from the active loop prominences which are associated with solar flares. It is therefore unlikely that a single prominence theory would cover all the observed features.

The loop prominences originate in regions of strong magnetic fields and the quiescent prominences in regions of very weak magnetic field.

A common element in the theory of formation of prominences is that they are supposed to be caused by a thermal instability in the corona leading to cooling and associated compression of the coronal matter. For quiescent prominences we find almost pressure equilibrium between the prominence and the ambient corona while for active prominences like loops the pressure has increased and moreover it seems that the magnetic field strength is so high that $\beta \ll 1$ in active region prominences.

It is therefore expected that active region prominences can only be formed out of coronal material if a sufficiently strong compression occurs. It has been shown by Kleczek (1958) that a compression first causes a strong heating of the medium after which a rapid cooling occurs if the density becomes so high that the radiative losses are larger than the compressional heating. A recent calculation made by de Bibhas (1973) leads to a formation time $\tau \approx 10^3$ s for loop prominences originating out of coronal plasma density $n_e = 10^9\,\mathrm{cm}^{-3}$ by constriction of a tube of force due to an axial current.

The fact that loop prominences become first visible at the top is consistent with this theory since the compression in the top of a field loop is slightly higher than at the bottom. Moreover, the observation that loop formation is associated with the appearance of the yellow coronal line confirms that a coronal loop is first heated when it is strongly compressed.

However, there are many more types of active region prominences and it seems that not all types can be explained by a simple compression of coronal matter. It appears that for some prominences the mass balance requires an additional injection of material from below. For a comprehensive discussion of all the mechanisms that have been recently suggested to explain the active prominences the reader is referred to the recent monograph on prominences by Tandberg-Hanssen (1974).

The second class of prominences which in general have a much longer lifetime and appear as long filaments on the disk when observed in Hα are the quiescent prominences. These prominences are not only quiet in the sense that they seem to be long lasting with lifetimes of the order of several weeks to sometimes several months, but they also appear above those regions of the Sun which do not show much activity. When they occur in active regions they do so when there are no rapidly varying or complex magnetic field structures. If the magnetic field in a centre of activity starts to change the quiescent filaments dissolve rapidly.

The most elongated and best developed quiescent prominences occur in between active regions at the boundaries where the fields of both regions have expanded and met each other, or in regions of very weak fields in between remnants of old centres of activity. The formation takes several days and seems to occur at greater heights in the corona than for the filaments in active regions (Martin 1973). The fact that prominences occur predominantly at the base of coronal streamers led Kuperus & Tandberg-Hanssen (1967) to the suggestion that they originate in magnetically neutral sheets in the corona. This has recently been confirmed by Mercier (1973), who found a strong correlation between type III radiobursts which presumably occur along open streamers (McLean 1970) and the existence of filaments. It has been shown by Raadu & Kuperus (1973) that in the lower parts of the corona linetying of magnetic field prevents the coronal matter from being compressed, while in the higher coronal layers the cooling proceeds much too slow. Therefore there is an optimal height for the formation of prominences. The 'condensation' of matter in a prominence is associated with the motion of the ambient corona. As long as the velocity is smaller than the minimum value of the sound velocity and the Alfvén velocity the condensation occurs almost at pressure equilibrium, since at large distances from the sheet matter and the field move together; the matter piles up in the sheet. If we adopt for the corona $n = 10^8\,\mathrm{cm}^{-3}$, $T_0 = 1.6 \times 10^6\,\mathrm{K}$ and $B = 5 \times 10^{-4}\,\mathrm{T}$ total pressure equilibrium throughout the sheet requires a gas pressure enhancement in the sheet $p_1/p_0 = 100$, where p_1 is the pressure in the sheet. Owing to the associated density enhancement strong cooling results. Neglecting heat conduction, the dense plasma in the sheet cools within a cooling time given by

$$\tau_c \approx \gamma p/(\gamma - 1)\, q_r, \tag{6}$$

where q_{rad} is given by equation (4) (Raadu & Kuperus 1973). The lower dense parts cool faster than the upper regions, thus causing a temperature gradient along which heat can flow into the sheet from above unless the topology of the field changes.

Changes in the field topology in a time much shorter than the diffusion time can occur in a neutral sheet due to the tearing mode instability (Furth, Killeen & Rosenbluth 1963). Owing to this instability a current sheet is first split up into many isolated currents, which merge into one current filament at a rate a few times the linear growth rate $\tau_{\mathrm{t.m.}}$ (Dickmann, Morse & Nielson

1969). The growth rate for the linear tearing instability is approximately given by

$$\tau_{\mathrm{t.m.}} \approx \tau_{\mathrm{A}}^{\frac{1}{2}} \tau_{\mathrm{d}}^{\frac{1}{2}}, \tag{7}$$

where $\tau_{\mathrm{A}} = l/C_{\mathrm{A}}$ and $\tau_{\mathrm{d}} = 4\pi\sigma l^2/c^2$ are the Alfvén time and the diffusion time, c_{A} is the Alfvén velocity, l the thickness of the current sheet and σ the electrical conductivity. Comparing equation (6) and equation (7) we find that topology changes occur in about the same time necessary for the formation of a cool filament if the typical length scale $l \approx 200$–$300\,\mathrm{km}$ (Kuperus & Tandberg-Hanssen 1967). Once the instability has set in the plasma may cool further since it is now thermally insulated from the hot surrounding corona due to the strong reduction of the thermal conductivity in directions perpendicular to the magnetic field. Owing to the pressure of the ambient plasma and field one would expect the sheet to collapse to still higher densities. Observations show that there are no great differences in gas pressure between a prominence and its surroundings. Presumably the tearing instability has prevented the sheet from this collapse (Kuperus 1974).

Once the current sheet has contracted into one cool dense current filament it tends to sink under its own weight. Previous theories invoked depressed horizontal fields to support the prominence (Kippenhahn & Schlüter 1957). However, it can be shown that a current filament embedded in a neutral sheet, while the surrounding magnetic field lines are tied to the photosphere is subject to a levitating force which is inversely proportional to the height above the photosphere (Kuperus & Raadu 1974). This force, which is caused by currents induced in the photosphere where the field lines are 'anchored', should be balanced by the weight of the prominence. Thus quiescent prominences can be formed and maintained at certain heights in the corona. They indicate the border lines of regions of opposite polarity where a magnetically neutral sheet is present which extends far outward as a coronal streamer. Quiescent prominences manifest a process of slow magnetic field reconnexion in the corona.

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Discussion

Dr A. H. Gabriel

I was very interested in your prominence model. I would suggest that such a mechanism might apply also to the formation of spicules, if one rejects the usual assumption that spicule material moves upwards at high velocity. Spicules would then be formed by the condensation of coronal material along a neutral sheet. The apparent growth rate of the spicule does not then represent the velocity of the material, but only the velocity of advance of the luminosity front. This then eliminates problems associated with the requirements for a large energy source or large mass transport. I suggest that this is not inconsistent with the observations. There is strong evidence for large Doppler widths, which would in this model be due to the ions retaining their coronal thermal energies. The evidence for Doppler *shifts* is however much less certain.

M. Kuperus

Your suggestion is very interesting and has been made before by several other people (e.g. Pikelner, Uchida). However, I do not think that one can reject the fact that spicules consist of upward moving 'cool' material as has been derived from the observed Doppler shifts. The condensation of coronal material thus has to be associated with upward motion.

A possible mechanism along these lines has recently been put forward by Glengross, who suggests that the cooling occurs like prominences in neutral sheets, which originate when twisted braids of field lines emerge. The motions result from the expected unwinding of the field when rising in the layers with a lower pressure.