A STUDY OF SPECTROSCOPIC AND KINEMATIC CHARACTERISTICS OF PECULIAR A-TYPE STARS

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In Memory of

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ABSTRACT

A spectroscopic survey of the different types of peculiar-A (Ap) stars has been made to discover the systematics of behavior of the various elements. Spectroscopic results are preceded by an account of some of the "gross Properties" investigated from the published data.

Gross Properties: (Chapter II): Under this group of investigations (a) intrinsic frequency distribution of the various peculiarity types is derived, (b) the controversy "whether Ap stars are too blue for their HD spectral types" is resolved, (c) mean absolute magnitudes of the different peculiarity types are derived by a statistical analysis of trigonometric parallaxes, and (d) space velocities are calculated which show significantly different solar motions for the different groups, but the velocity dispersions are found to be comparable.

Spectroscopic Studies: This study is based mainly on 21 to 45 (A/mm) dispersion plates, in the UV and blue regions of the spectrum of 40 Ap and 15 normal stars. k (A) (Chapter III) By eliminating the effects of temperature, electron pressure, continuous absorption and microturbulence, "strengthening" and "weakening" of various elements has been estimated. From these estimates (Chapter VI): (i) behavior of the various elements in Ap

stars is described, (ii) several interesting correlations between the observed anomalies of various elements have been discovered, and (iii) for Ap stars, a new system of classification is proposed which is based on the general pattern of abnormalities rather than on single most prominent features. (B) Extensive tables of reliable blend-free UV and blue lines are given for detection of the various elements (Chapter III). A similar table is given for the detection of all the rare-earths (Chapter V). In different stars, relative distribution of the rare-earths is found to be different. (C) A survey of BeII $\lambda\lambda$3130-31 is made. In seven Ap stars, from "Mn" and "Si-4200" groups, this doublet is very strong. A quantitative Be abundance analysis is given (Chapter IV).

Various explanations for the appearance of the abnormal line strengths - including the predicted nuclear processes - are discussed in the light of the present results. Some specific suggestions for further work are also made (Chapter VI).

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CHAPTER I

Section I-A

A. Background for the Peculiar-A Star Studies

During the past six decades many authors have studied the peculiar A-type stars (Ap stars) and the spectrum variables; however, understanding of these objects is still not in a satisfactory state. Many of the earlier studies were based on rather low dispersion spectrograms on which only the most conspicuous features could be seen. The 2.8 magnitude star α^2 CVn, of course, has most often been studied, often with high dispersion. Struve and Swings (72) believed that explanations for abnormal line strengths which involved abundance anomalies were "escapist ideas". But more recent work (Searle and Sargent (69)) shows, with reasonable certainty, that at least some of the abundance anomalies are real.

Efforts have been made to explain these abundance anomalies by processes of nucleosynthesis which are assumed to have occurred under abnormal conditions on the surface or in the upper layers of the stars. In this respect, however, there has been a tendency to consider all the Ap stars as one group ignoring the diversity in the behavior of various elements in the different peculiarity groups. Such lumping together of peculiarities can present an unreal picture, making it difficult to explain them by the processes of nucleosynthesis or otherwise. It is, of course, known that abundances vary in the different peculiarity types; but the number of Ap stars for which abundance analyses have been made so far is inadequate for delineating peculiarity patterns in these different peculiarity groups.

There is dispersion within each group, too. For example, as Searle and Sargent (69) point out, in some respects, κ Cnc is an atypical Mn-star. Consequently, if we observe only one or two stars from each peculiarity group, we cannot be sure which of the peculiarities are typical for that group and which are incidental to that particular star only. Obviously, for finding general patterns, it is necessary to observe a sufficiently large number of stars from each group. In his review article, Sargent (65) has collected results of abundance analyses (made by various observers over the past ten years) for eight Ap stars encompassing all the five peculiarity types. Therefore, it should not be surprising if some of these results do not depict a typical pattern for the whole group.

Usually it is tempting to think of the spectrum variable and the Ap stars as synonymous, in so far as their intrinsic physical structure is concerned - Deutsch (29) points out, "no criterion is known for concluding, from a single spectrogram of a peculiar A star, whether or not it will prove to be a spectrum variable." It is, however, important to realize that by virtue of the additional temporal dimension, spectrum variables provide a whole set of different possibilities for learning about the cause of the peculiarities. On the other hand, there are only a few spectrum variables suitable for extensive investigation.

Basically, two types of approaches are needed for further work on the Ap-stars and the spectrum variables. One of these, individual star studies, requires a concerted program of studying time variations of the various characteristics, such as radial velocities, line profiles, magnetic fields, etc. for each element in a single spectrum variable. Studies by Struve and Swings (73), Babcock (7), G. and M. Burbidge (20), and by others, on α^2 CVn, provide part of the results needed for such a study (unfortunately, α^2 CVn is not the most suitable star for a very detailed study.).

The other approach, that of statistical study, which will be used here, lies in studying a <u>large number</u> of different Ap-stars (temporal variations, if any, are ignored by considering each phase as a different star). Obviously, the type of information that is possible to obtain by studying temporal variations in a star, cannot be obtained by this type of study. However, it is important for the following reasons:

(A) Periodic spectrum variables have not been found among all the peculiarity types; well established spectrum variables are particularly lacking among the Mn stars. This in turn would imply that either,

(i) Mn-stars represent an unobservable aspect of the variable stars, or

(ii) Mn-stars are totally different from the peculiarity types that have spectrum variation. Furthermore, it _is_ possible that difference between spectrum variables and non-variables may be more than that of orientation.

(B) If correlations between abnormalities of some elements do exist, these can be discovered only by observing a large number of stars. Similarly, other types of correlations relating peculiarity types, magnetic field, line width, etc., can be discovered only by a statistical study of a large number of Ap stars.

Statistical study of the following characteristics and their correlations is particularly useful:

(i) Relative abundance of elements.

(ii) Line widths, rotational velocity.

(iii) Polarity and strength of magnetic field.

(iv) Color, absolute magnitude, peculiarity types, temperature, and electron pressure.

(v) Space distribution and space motions.

Naturally, a statistical study should include all the peculiarity groups.

Since, in this study, we shall be dealing with the Ap stars of the different peculiarity types, the following remarks are in order.

(a) Discovery of peculiarities depends upon dispersion used. For example, κ Cnc, which looks normal with an objective prism (HD spectral type is B8) is a very peculiar star at high dispersion.

(b) Though the number of known Ap stars is quite large, discovery of the different peculiarity types is very uneven. For example, strengthening of SiII and SrII lines can be detected with an objective prism and, therefore, a large number of Ap stars with Si and Sr peculiarities are known. On the other hand, lines of MnII, in the blue region, are very weak even at high dispersions and are seen only when Mn is overabundant by a large factor - strong lines of MnII lie in the ultraviolet region 3440-3500 Å . Consequently, relatively fewer Mn stars are known for $M_V > 5.5$.

(c) All the peculiarity classifications that have been made so far, are based on <u>apparent</u> strengthening or weakening of a single spectroscopic feature or of lines due to a single element. Consequently, differences in effective temperature and in degree of abundance abnormality of the various elements, result in a large number of peculiarity groups. The spectroscopic work in the present study enables us to group the Ap stars into a smaller number of groups based on patterns of intrinsic behavior of several elements.

(d) For investigations in Chapter II the peculiarity classification has been taken from the published sources. Though different criteria have been used at different times and with different dispersion used for classification of the Ap stars, the following criteria used by Miss Walther (77) are representative, in the mean, for the usual peculiarity types.

"'Cr' means that $\lambda 4077$ and $\lambda 4171$ of CrII are strong.
'Sr' means that $\lambda 4077$ aand $\lambda 4215$ of SrII are strong.
'Eu' means that EuII is implied by the fact that the general features of the spectrum match a typical chromium-europium star.
'4200' means that this unidentified line is strong.
'Si' indicates that the blend at $\lambda 4128$-31 is strong and that the general features of the spectrum correspond to a typical silicon star."

It is to be noted that according to these criteria, presence of EuII is inferred without actually observing any Eu lines. Furthermore, it should not be surprising then, if some of the earlier classifications are found to be completely wrong. For example, the sharp line star HD204411, listed as 'Eu-Cr' has EuII lines hardly detectable even on high dispersion plates (4.5 A/mm). Likewise, HD9996 listed as an "Si,Sr" star has actually weak SiII lines.

Section I-B

Plan of the Present Study

The present study is, then, a statistical investigation of a large number of stars. Investigations were made on many aspects of the Ap stars, but only those investigations which yielded positive and interesting results are reported here.

These studies can be classified into two groups, for which samples of stars used were different. For the first group of studies, observational information was abstracted from various published sources, and various "gross" properties, such as statistics, distribution, color, luminosity, space motion, etc. have been studied. Second, and the major part, covers spectroscopic studies on more than 40 Ap stars with about 10 normal stars used for comparison. More than two-thirds of the spectrograms used were taken for this program with the coudé spectrograph of the 100-inch telescope. The rest of the spectrograms were collected from the plate file of the Mount Wilson and Palomar Observatories. In addition to results on (a) "abundances", (b) Beryllium survey, and (c) Rare-earth survey, detailed tables for identification work have been presented which, it is believed, will be useful to future investigators undertaking extensive surveys. Emphasis in these tables has been given on establishment of criteria which are safe and blend free under the variety of circumstances encountered among the Ap stars and which provide quick estimates of strengthening or weakening of various elements in these stars.

The final section deals with discussion of various results and of various explanations for the Ap stars in the light of new observational results. Suggestions for further work are also made.

CHAPTER II

GROSS PROPERTIES

Section II-A

Relative Frequency of Peculiarity Types

For a statistical study of the Ap stars, a list was compiled from the following sources: (i) all the Ap stars in all the tables of the Babcock Catalogue, (ii) Bidelman's list of 124 stars brighter than m_V=7.0 (Jascheck and Jascheck (42), (iii) Ap stars brighter than m_V=6.8 in Bertand's Catalogues (11, 12), (iv) Eggen's list of bright A-stars (Eggen (32), and (v) two Mn stars, 53 Tau[1] and 112 Her, of recent discovery. From this compilation, stars now known to be Am stars, A-type stars with other types of peculiarities such as shell stars Ba-stars, faint-line stars, K-line stars, etc. were omitted.

The total number of Ap stars, N brighter than a given apparent magnitude, m., is plotted on a logarithmic scale in Figure II-1 in which the slope of the straight line is

$$\frac{\Delta \log N}{\Delta m} = .6$$

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[1] In Babcock's (4) and Bertand's Catalogs, HD 23193 is wrongly labeled as 53 Tau.

Up to $m_v = 5.5$ (N=74) this straight line relation is followed fairly well. (Though, in total, this agreement is good, it is possible, as we have seen in Chapter I, that some of the members might be spurious, while some genuine Ap stars, particularly Mn stars, might still be missing.) It is clear, nevertheless, that even if we leave room for such uncertainties, their effect must be minor; and we can draw reliable information on <u>intrinsic</u> distribution of the different types of peculiarities from this group of 74 stars with $m_v \lesssim 5.5$.

It has been pointed out in the previous chapter that on the blue spectrograms, Mn-stars need a relatively higher dispersion (preferably better than 40 A/mm) for their discovery while members of other groups can be discovered on low dispersion or objective prism plates. Therefore, in this group of 74 stars, we suspect only the Mn-stars to be significantly underestimated. An <u>estimate</u> of the incompleteness of Mn-star discoveries can be made by making a plot (Figure II-2) for the known Mn-stars in the fashion of Figure II-1. Figure II-2 indicates that we should have about 20 Mn-stars brighter than $m_v = 5.5$ while the number of known Mn-stars is 14.

In Table II-1, this "uncorrected" figure for the Mn-stars is given within parentheses. In this table, for each peculiarity group, number of stars, per 10^6 pc^3, and their percentages are also given. Two sets of values are given: For (A), the calculations are based on the mean absolute magnitudes derived in the present work (see Table II-4a), and for (B), the calculations are based on the mean B-V value for each group and Eggen's (32) relation, $M_V = 7.8\ (B-V) + .^m35$.

Thus, for the first time, we have obtained an intrinsic frequency distribution for the different groups of the Ap-stars. The peculiarity classification for all the stars of this study have been taken from published sources. (In a recent paper by Osawa (62), which was kindly brought to my attention by Dr. Deutsch, classifications are available for all but 13 of these 74 stars. Osawa's classifications for these stars are in good agreement with the previous ones.) The group "Am-like" contains stars with $+.2 \leqslant B-V \leqslant +.3$, and is discussed in Section II-C.

It should be stated that an implicit assumption in the relation, $\frac{\Delta \log N}{\Delta m} = .6$ is that the different groups have radially uniform distribution in space around the sun..

For a statistical discussion, these 74 stars will be used in the ensuing sections of this chapter.

Section II-B

Colors

The (B-V) colors and V magnitudes for these stars are given in Table II-5. The values adopted here, are often a mean of different values published. Following are the sources for colors: Abt and Golson (1), Eggen (32), Crawford (24), Stoy and Cousins (71), Osawa (61), and Hoffleit (40).

(U-B) Vs. (B-V): Provin (64) found that the Ap stars he observed lie along the main-sequence stars on the two-color diagram. Later, Abt and Golson found this to be the case also with the sharper line Ap stars in which Babcock had measured the magnetic field. However, in view of the suggestions (Bidelman (13)) that Ap stars might be evolved stars returning to the main sequence on the color-luminosity and two-color diagrams, it becomes interesting to re-examine how close to the main sequence Ap stars lie. Figure II-4 is the two-color diagram for 49 of the 74 stars. For others, one or both colors were not available. These include both the sharp line and the broad line stars.

We notice a very curious fact in this diagram - all the typical Mn stars have observed B-V colors in an extremely narrow range (B-V = -.10 ±.01). Of the 12 Mn stars plotted in this diagram, ten fall within this range.

Of the other two, ι CrB is only a mild example of the Mn group, while ϕ Her is atypical because it has conspicuously large overabundance of Cr; in other Mn stars, Cr is underabundant or at the most, normal. Sargent (65) gives a somewhat wider range, B-V = -.05 to -.10, for spread in colors of the Mn group.

These Ap stars lie remarkably close to the main-sequence on the two-color diagram. A similar (U-B) - (B-V) plot was made for the stars in Table 1 of Morgan's paper (58). Scatter for the Ap stars was not at all greater than that for the normal main-sequence stars. Outstanding departure is found in the case of 21 Per which falls at the same position as the AOIb star η Leo (B-V = -.01; U-B = -.24). But, the Balmer confluence point in 21 Per has a value appropriate for a main-sequence star. In view of its bright apparent magnitude, $5^{m}.01$, and $\ell^{II} = 152°$, $b^{II} = -24°$, reddening by interstellar absorption does not seem to be likely. Therefore, energy distribution scans of 21 Per should be made for determining the reason for its abnormal colors. It might be added that there are indications of variations in radial velocity of 21 Per (Hoffleit (40)), and consequently, it could be a spectroscopic binary.

Distribution of B-V Colors: Distribution of Ap stars of Table II-5, as a function of B-V, is shown in a histogram in Figure II-5(b); total number of stars being 66. The most interesting feature in this distribution is an absence of Ap stars in the range $+.12 < B-V \leq +23$; then again we have four Ap stars in a very narrow range $B-V = +.24$ to $+.27$. Is this paucity of the Ap stars due to intrinsically fewer A-type stars in general in this color range, or is this due to a rapid decrease in the proportion of the Ap stars at $B-V \simeq +.12$? To answer this question, a count was made from Eggen's list (32) which contains A stars satisfying the conditions (a) $M_V \leq 5.5$, and (b) HD spectral type B8 or later but $(B-V) < +.30$. In the present count only the luminosity class V stars within the range $+.10 \leq B-V \leq +.30$ were included. Figure II-5 (c) shows the distribution for normal stars. A similar count for Am stars is shown in Figure II-5 (a). None of the Figures II-5 (a) and (c) gives an indication of any paucity of stars for $+.12 < B-V < +.23$. According to the above mentioned count, there are 75 normal stars in this range, and since Ap stars comprise 10-13% of the normal stars, we should expect at least eight Ap stars in this range which we do not find. This means the second alternative, that the proportion of the Ap stars is very low for $B-V > +.12$, is correct. But then why, suddenly, are there 4 Ap stars within a very narrow range of B-V? It seems that these stars, intrinsically, belong to region B-V $\simeq$ +.10 but are moved red-ward due to line-blanketing effects. (Baschek and Oke (8) have determined blanketing correction for β CrB, and on deblanketing, this star moves to B-V = +.1. On 10A/mm plates of the sharp line Ap star, γ Equ, FeI to FeII line ratio indicates a spectral type about A4, which means B-V $\simeq$ +.12.) But then one asks <u>why blanketing effect does not spread stars red-wards in a continuous fashion</u>, and why there is an isolated group of four stars within a small range of B-V? It is possible that these four Ap stars represent a different group, which we might call "Am-like". The existence of a small group of Ap stars around B-V $\sim$ +.25 has been recognized by C and M. Jaschek (42), but for different reasons.

Mean colors for the different groups of Ap stars are given in Table II-2. Blanketing corrections are based on the correction determined by Baschek and Oke for 52 Her(Sr) and β CrB ("Am-like").

<u>(B-V)-HD Spectral Type</u>: The original finding by Deutsch (27), that Ap stars are too blue for their HD spectral types, has been a subject of controversy, and even in the recent literature diverse views have been expressed. For example, Jaschek and Jaschek (43) have maintained

that only the "Si-4200" stars are too blue for their hydrogen spectral types. Reliable B-V colors are now available for a large number of Ap stars and it should be a simple matter to reach a definitive conclusion. We can trace the cause of this controversy, and find that it lies in an over-emphasis, in recent years, on the (B-V)-peculiarity type correlation.

Both "Mn" and "Si-4200" groups contain hotter Ap stars, but constitute two <u>very</u> distinct groups. On the other hand, (genuine) Ap stars of Si-4200, Si, and Si-Cr-Eu, groups of the earlier classifications, contain several common characteristics, and their differences in spectral appearance are largely due to temperature differences. (See Chapter VI-C and Figure VI-16.) Once this fact has been recognized, we should not any more insist on arranging the different peculiarity groups in order of their mean B-V colors. Specifically, we should keep the Mn-group separate, and we should not put it betwen the "Si-4200" and "Si-Cr" groups as has been done by earlier investigators.

Table II-3 gives values as read from Figure 2(b) of the paper by the Jascheks (43). "Blanketing corrections" are interpolated values between the corrections for 52 Her (B-V = +.08) and β CrB (B-V = +2.7), derived by Baschek and Oke (8). Examination of Table II-3 leads us to state <u>Deutsch's original conclusion in the following modified form</u>:

Except for the Mn-stars (incidentally, most of the Mn-stars were not recognized as "peculiar" in the HD Catalog), other Ap stars - especially those of the Si groups - have colors which are too blue for their HD spectral types. This color-anomaly is greatest for the hottest Si-stars.

(B-V)-Temperature Relation: The temperature-(B-V) relation, according to Oke (60), for the normal stars is

$$\theta_e = (B-V) + .53 \quad .$$

Searle and Sargent (69) find the relation

$$\theta_{ionization} = (B-V) + .50$$

applicable to the normal and Ap stars. (Except for the Mn-star 53 Tau, the conclusion of Searle and Sargent about the Ap stars is based on temperatures derived using lines formed in the deeper layers of stellar atmosphere. For the limited objective of the present study we shall derive θ values from B-V colors using the above relation. But it should be mentioned that for the reasons given below we suspect that the validity of this θ-(B-V) relation might be limited only to the high excitation lines formed in the deeper layers of the atmosphere which are represented by the continuum radiation.

In Figure II-6, values of θ_{ion} (obtained (i) from spectroscopic analyses of lines formed at relatively upper layers of the atmosphere and (ii) from the B-V colors), as listed in Table II of Sargent (65), are plotted for the stars of the usual peculiarity types. Crosses represent β CrB and γ Equ after applying blanketing correction to B-V derived for β CrB by Baschek and Oke (8). In this figure, disagreement between the two values of Θ_{ion} derived from B-V colors and spectroscopic analysis is very conspicuous except for the Mn-stars. This disagreement gives rise to the doubt about the validity of the above relation for use with lines formed in the upper layers of the atmosphere. Though both Mn-stars and Si-stars are among the hotter Ap's, these discrepancies are large for the latter only. In this connection, it may be recalled that in Si-stars, on the average, magnetic fields are much stronger than in the Mn-stars, and it is possible that these discrepancies in Si-stars are caused by a steeper temperature gradient in the atmosphere resulting from these strong magnetic fields.

Section II-C

Absolute Magnitudes

Absolute magnitudes for some cooler Ap stars were derived by Deutsch (27) and Eggen (30) using trigonometric parallaxes. Subsequently, C. and M. Jaschek (42) calculated mean absolute magnitudes from proper motions, for the different peculiarity groups of the Ap stars. They used the same value of solar motion, 20km/sec, for all the groups. In actuality, we find a considerable difference between the solar motion for the cooler and the hotter Ap stars (Table II-6). Evidently a new calculation became necessary. But, if we determine components of the solar motion and mean absolute magnitude, both from the proper motions, the accuracy of the results would be very low. In the present section, therefore, we shall calculate mean absolute magnitudes statistically from a large number of trigonometric parallaxes.

An outline of the procedure used for these calculations is as follows:

(1) For each group, mean absolute magnitude $\bar{M}$ was calculated from the formula:

$$10^{0.2\bar{M}-1} = \frac{\sum_i p_i\, 10^{-0.2 m_i}}{\sum_i 10^{-0.4 m_i}}$$

where p is trigonometric parallax, and m is apparent magnitude. Since errors in p have a normal distribution, this formula permits the use of all the stars with known trigonometric parallax (57 stars); this in turn mostly cancels out the cut-off error (normal error distribution in parallax produces a skew distribution in absolute magnitude value). Earlier investigators, who used trigonometric parallaxes, calculated absolute magnitudes individually for each star, and therefore, could use only a small number of Ap stars - mainly the cooler ones - which have large trigonometric parallaxes.

(2) We also want to compare the luminosity of the Ap stars with that of the normal stars; it is most desirable that the absolute magnitude of the latter be determined in an identical manner so that systematic errors are eliminated. Therefore, all the luminosity class V stars, brighter than $M_V = 4.0$ and with trigonometric parallaxes available, were selected from Eggen's list (32). These stars, arranged in a sequence of B-V colors, were divided into groups such that no group contained too few stars for the purpose of taking mean. For a few of the brightest normal stars, absolute magnitude was calculated separately.

(3) The resulting mean B-V colors and absolute magnitudes for different groups of Ap and normal stars are given in Table II-4a. For comparison, results of C. and M. Jaschek are also given, in Table II-4b.

On the basis of the above calculations, we note these points:

(a) Eggen's latest (32) relation, $M_V = 7.8$ (B-V) $+^{m}.35$ gives absolute magnitudes which are on the fainter side for B-V $<$ -.05 (Consequently, the distance modulii are too small. We shall refer to this point in Section II-E.)

(b) Values of mean absolute magnitude for the Ap stars, particularly the hotter ones derived in the present work, are more luminous than the values derived by the Jasheks.

(c) No significant difference between the absolute magnitudes of the field Ap stars and the field normal stars of the same color is found.

<u>Section II-D</u>

<u>Space Motions</u>

Space velocities, with respect to the sun, for stars of Table II-5 have been calculated. Proper motions and radial velocities have been taken from the Bright Star Catalogue (40). Calculations have been made for two sets of distance values: (a) for distances obtained by using the color-absolute magnitude relation derived in the previous section (Table II-4a). (b) for the distance modulus values given by Eggen (32) (in this second set, for those stars for which Eggen has not given distance modulus, Eggen's relation $M_V = 7.8 (B-V) + .35$ was used; the distance modulii thus derived are given in parentheses in Table II-5). As we have pointed out earlier, Eggen's relation gives absolute magnitudes which are too faint for $B-V < -.05$ and consequently gives small space motions. The differences in space motion values derived from (a) and (b), therefore, will give an upper limit to errors in space motion, on account of uncertainty in distances.

The U-, V-, and W- axes are pointed in the directions $\ell^{II}=180°$, $b^{II}=0°$, $\ell^{II}=90°$, $b^{II}=0°$; and $b^{II}=+90°$ respectively. Formulae of components, U, V, and W, of space motion of stars, relative to the sun, are given in Appendix I. Space motion for a group of stars is given by

$$S \equiv (\bar{U}^2 + \bar{V}^2 + \bar{W}^2)^{1/2} ,$$

dispersions σ_u, σ_v, and σ_w, are given by

$$\sigma_u = \left[\frac{1}{n} \sum_{i=1}^{n} (U_i - \bar{U})^2 \right]^{1/2} , \quad \text{etc.,}$$

and $$\sigma = (\sigma_U^2 + \sigma_V^2 + \sigma_W^2)^{1/2}$$

Components of space motion for 66 Ap stars are given in Table II-5; and values of space motions for the different peculiarity groups, dispersions, etc. are given in Table II-6. Values for the normal stars are calculated from data given by Allen (2). We make the following conclusions from these results:

(i) The different peculiarity groups do not possess a common mean space motion. Departure for the "Sr" group is particularly noteworthy.

(ii) Dispersions for the different groups are not significantly different, and therefore, the interpretation that the bluer Ap stars are kinematically more homogeneous and cooler ones are heterogenous (Eggen (33)), is not supported by these calculations.

(iii) It is very interesting to note that ten Mn-stars, out of the 14 known Mn-stars with M_V 5.5, lie within a small region in the UV plane and with $(\bar{U}, \bar{V}, \bar{W}, \sigma) =$ (+13, -31, -6, 11). Some of these have already been found to be members of the Pleiades "group" by Eggen (32). But, in the Pleiades cluster itself, no Ap stars have been discovered with certainty - spectral classification for only the brighter members (mostly earlier than B8) of the Pleiades cluster are given by Johnson and Morgan (44). It will be valuable, therefore, to investigate if the Pleiades cluster itself contains Ap stars, and more particularly Mn-stars.*

The other Mn-stars, particularly υ Her and π'Boo, have very different space velocities so that a combined value for dispersion, σ, becomes large, as for the other groups. Therefore, these other Mn-stars must belong to some different "group".

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*Dr. Deutsch points out that an Mn-star has been found in the Pleiades cluster.

Peculiar A-type Stars in the Galactic Clusters

Data on occurrence of Ap stars in galactic clusters has been compiled by C. and M. Jaschek (43) and by Meadows (49). After examination of all the original sources of information cited in these papers, the author has the view that in several cases further observations are necessary to make those spectral classifications secure. Moreover, such investigations need to be made much more extensively before a reliable interpretation of those observations can be made. For example, in younger clusters, spectral classifications have been made mainly for stars at the most luminous end of the sequence, and only for very few stars down to a level where cooler Ap's and Am's are likely to be found. Therefore, the contention that the development of peculiarities follow a time sequence cannot be taken as established. In fact there is a possibility that the suggestion of such an evolutionary sequence might be a result of observational selection.

Eggen's work on stellar "groups" does not provide evidence in favor of the hypothesis that Ap stars are the most luminous members, in a cluster or group, evolving off the main sequence. See, for example, Figure 2 in Eggen (31), and position of HR 710, (M_V =+2.29), a well-known Sr-star, well below the "break-away" point, (M_V = +0.5), in the color-magnitude diagram for the "Sirius Group" (Eggen (31)).

TABLE II-1

Peculiarity Group	No. of Stars to M_V=5.5	No. of Stars/ $10^6 pc^3$		Percent for a Given Volume	
		A	B	A	B
Mn	20(14)*	.61	1.6	7	13
Si-4200	12	.26	.8	3	6
Si-Cr	20	1.6	2.0	19	16
Cr-Eu	12	3.3	2.6	39	21
Sr	12	2.7	5.5	32	44
Total	70	8.5	12.5	100	100
"Am-like"	4	3.5	13	41	102

* "Uncorrected" value for the number of Mn-stars is given in parentheses.

A = These calculations are based on the mean absolute magnitudes derived in the present work (See Table II-4).

B = These calculations are based on the mean B-V value for each group and Eggen's relation $M_V = 7.8(B-V) + {}^{m}\!.35$

TABLE II-2

Group	Number of Stars	Mean B-V Color	Blanketing Correction
Mn	14	-.10	0
Si-4200	12	-.11	0
Si-Cr	16	-.08	0
Cr-Eu	11	-.00	0
Sr	10	+.06	-.05
"Am-like"	3	+.26	-.15

TABLE II-3

Group	Mean HD Sp. Type	Mean B-V	Color Anomaly	Blanketing Correction
Mn	B8-B9	-.09	-.01	.0
Si-4200	A0	-.13	-.13	.0
Si	A0	-.06	-.06	.0
Cr-Eu	A1	+.02	-.03	.0
Cr-Eu-Sr	A2	+.07	-.01	-.03
Sr	A3	+.13	+.04	-.07

TABLE II-4a

Absolute Magnitudes for Normal and Ap Stars

	Ap Stars			Normal Main Sequence Stars		
Group	Mean B-V	n	$\overline{M}_V$	Mean B-V	n	$\overline{M}_V$
Mn	-.08	6	-1.0			
Si-4200	-.10	10	-1.2	-.04	14	-0.2
Si-Cr	-.06	17	-0.3	+.00	11	+0.2
Cr-Eu	+.01	10	+0.6	+.05	10	+0.7
Sr	+.08	10	+0.5	+.09	7	+1.6
"Am-like"	+.26	3	+1.4	+.16	8	+1.6

TABLE II-4b

Absolute Magnitudes From Jaschek and Jaschek (42)

Group	"B-V"	$M_{Correct}$
O	-.13	+0.1
A	-.09	-0.6
B	-.08	-0.1
BC	-.07	
C	+.01	+0.6
CD	+.09	+1.0
D	+.13	+1.8

TABLE II-5

SPACE VELOCITIES FOR THE INDIVIDUAL AP STARS

				A				B			
HR	Type	B-V	V	m-M	U	V	W	m-M	U	V	W
7552	Cr	-.15	5.30	5.6	9	-6	-17	(5.6)	9	-6	-17
8151	Cr	+.02	4.81	4.5	19	-2	-22	3.35	11	-1	-14
5475	Mn	-.10	4.94:	6.0	-3	10	-4	5.0	-2	6	-3
5843	Sr	+.04	5.33	4.9	-15	6	-12	4.68	-14	5	-10
6234	Cr	-.02	5.24	5.2	3	-28	-8	5.05	4	-27	-8
6117	Cr	+.01	4.56	4.4	-15	-9	-22	4.10	-13	-8	-20
7287	Si	-.08	5.12	5.9	5	-2	-4	5.65	5	-2	-3
5747	Fp	+.27	3.69	2.4	29	-18	-0	2.20	27	-17	-2
8452	Si	-.12	5.46	6.8	28	8	-16	(6.05)	20	6	-12
5971	Mn	-.05	5.00	5.5	8	-27	-3	5.00	8	-23	-5
7113	Mn	-.09	5.20	6.1	3	-24	0	6.50:	1	-25	1
8097	Fp	+.27	4.64	3.3	-3	-34	-20	3.40	-3	-35	-21
9031	4200	-.15	5.17	7.0	18	-1	-16	(5.98)	11	0	-15
6485	Si?		4.52								
6997	Mn	-.11	5.39	6.6	4	-35	13	5.55	7	-30	5
7395	4200	-.11	5.12	6.3	13	-19	-1	5.85	12	-19	-1
6023	Mn	-.06	4.26	4.8	19	-14	-4	4.65	18	-14	-5
6254	Sr	+.08	4.81	4.2	-21	-3	-6	3.80	-17	-2	-5
5982	Mn	-.10	4.60	5.7	-54	7	-16	5.25	-44	6	-13
5857	Si	-.07	5.39	6.1	41	-36	14	5.70	34	-32	9
8911	Sr	+.03	4.90	4.6	11	-41	-26	4.30	10	-36	-22
7049	Si	-.12	5.05	6.3	19	-30	-4	5.50	13	-29	-6
8216	Cr	+.09	5.28	4.6	22	-14	-9	4.25	18	-14	-8
6920	4200	-.10	4.18	5.3	16	-21	-3	5.60	19	-22	-3
5187	Cr	-.09	5.49	6.4	8	-17	2	(5.85)	6	-14	1
7879	Sr	+.07	5.20	4.7	-0	11	-2	4.30	0	11	-1
5109	Si	-.04	5.52	5.7	5	-17	-3	4.35	2	-11	-5
5291	Si	-.05	3.64	3.9	10	-17	-10	3.65	9	-16	-11

TABLE II-5 Cont.

				A				B			
HR	Type	B-V	V	m-M	U	V	W	m-M	U	V	W
15	Mn	-.11	2.02	3.2	+8	-36	-26	2.80	+6	-32	-20
4915	Si	-.11	2.92	4.1	69::	-29:	-4	3.45	51::	-21:	-4
4905	Cr	-.02	1.76	1.8	-13	2	-8	1.60	-12	2	-8
707	Sr		4.51								
1643	Si		5.32								
546	Cr		4.8:								
4072	Si	-.06	4.93	5.4	6	-13	3	5.05	5	-11	3
873	4200	+.00	5.10	5.1	+6	-7	-16	5.30	+6	-8	-17
596	Sr		5.96:								
1971	Cr	+.03	5.47	5.2	-5	-0	-6	4.89	-5	-1	-6
1268	4200	-.12	5.19	6.5	3	-51::	-18	5.85	2	-38::	-13
1732	4200	-.17	5.36	7.5	34	-43::	-7	5.00	31	-11::	-3
1341	4200	-.14	5.24	6.9	16	-58::	-6	5.50	14	-30::	-5
2095	Si	-.08	2.69	3.5	31	-20:	4	2.95	30	-15:	4
1339	Mn	-.09	5.28	6.2	13	-43::	-6	4.85	11	-23::	-5
1791	Si?	-.13	1.65	3.1	8	-32:	-15	2.31	8	-22:	-11
2425	Cr	-.01	5.51	5.5	16	-10	-12	5.24	15	-9	-11
1638	4200	-.07	4.67	5.4	13	-24	-8	4.90	14	-19	-7
2033	Si	-.04	5.54	5.7	-5	2	-12	5.50	-5	2	-9
3595	Si	-.05	5.44	5.7	-12	-1	-11	5.50	-12	-1	-11
1100	Si	-.14	5.22	6.9	14	-28	2	6.00	12	-20	-2
1702	Mn	-.11	3.28	4.5	15	-30	-4	3.90	16	-26	-6
3623	Mn	-.11	5.24	6.4	26	-22	-6	6.00	24	-20	-3
1240	4200	-.12	4.66	6.0	22	-10	-11	5.35	+20	-10	-13
612	4200	-.16	4.68	6.7	15	-2	-16	5.50	+10	-3	-17
4752	Cr	-.05	5.30	5.6	6	-21	-5	4.55	4	-13	-4
2657	Mn	-.12	4.10	5.4	16	-26	-3	4.69	17	-25	-3
3500	Mn	-.10	5.19	6.3	24	-36	-14	4.80	21	-30	-0
4766	Sr	+.05	5.44	5.0	1	-9	-1	4.55	1	-7	-1
2727	Cr	+.06	5.30	4.8	-6	9	-11	4.48	-5	+8	-9
2746	Si	-.03	4.89	5.0	-40:	-5	-32:	4.65	-34:	-5	-28:

TABLE II-5 Cont.

HR	Type	B-V	V	A m-M	A U	A V	A W	B m-M	B U	B V	B W
1465	4200	-.10	3.26	4.4	7	-31	-4	3.65	5	-28	-8
2683	Si		5.30								
4327	Sr	+.03	5.14	4.8	44:	-13	-2	(4.56)	40:	-12	-2
4817	Mn	-.09	4.64	5.5	14	-36	-17	5.40	14	-34	-16
5303	Fp	(+.26)	4.90								
5463	Fp	+.24	3.18	1.9	18	-26:	-17:	1.05	11:	-19:	-12:
6204	Si	-.08	5.12	5.9	12	-11	-2	5.5	10	-9	-2
5105	Sr	+.04	4.93	4.5	-13	3	-18	4.30	-11	3	-17
5719	Si		5.38								
5523	Sr	+.06	5.38	4.9	17	-25	4	(4.56)	15	-22	3
5652	Si	-.09	4.53	5.4	15	-30	-14	5.30	15	-29	-13
8949	Sr	+.08	4.70	4.1	4	-7	-21	3.70	2	-6	-21
5313	4200	-.12	4.95	6.3	14	-51:	4	5.60	10	-37:	3
8937	Mn	-.10	4.36	5.5	46	-10	-18	5.00	37	-8	-15
6153	Sr	+.12	4.46	3.7	-3	8	3	3.15	-3	6	3

TABLE II-6

SOLAR MOTIONS FOR THE DIFFERENT GROUPS

		A									B			
Group	B-V	n	U	σ_u	V	σ_v	W	σ_w	σ	S	U	V	W	S
Mn	-.10	14	+11	9	-22	15	-8	8	19	26	+11	-20	-6	24
Si-4200	-.11	12	+15	8	-22	15	-8	7	18	27	+13	-17	-8	23
Si-Cr	-.08	16	+12	15	-16	13	-6	9	22	21	+10	-13	-6	18
Cr-Eu	-.00	11	+4	12	-9	11	-11	7	18	14	+3	-8	-10	13
Sr	+.06	10	+0	13	-5	12	-8	10	20	10	+0	-4	-7	9
Am-like	+.26	3	+14	13	-26	7	-11	9			+12	-25	-12	
BO	-.31		+11	13	-18	7	-8	4	16	22				
AO	.00		+10	12	-11	15	-6	7	20	16				
FO	+.30		+11	20	-12	20	-6	12	31	16				

Velocities and dispersions are in Km/sec.

n is the number of stars in each group used for computation.

A - calculations based on luminosities derived in the present work.

B - calculations based on Eggen's linear relation.

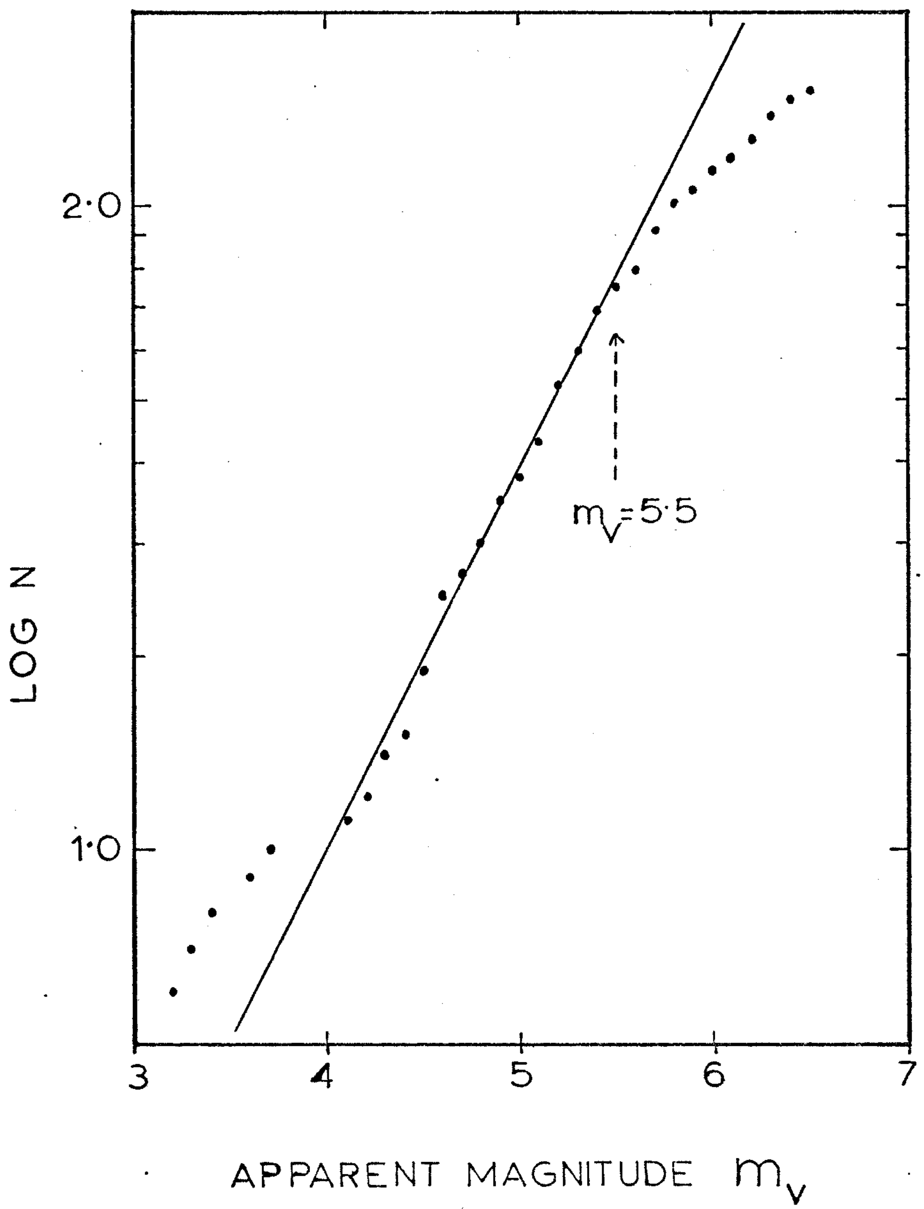


Figure II-1

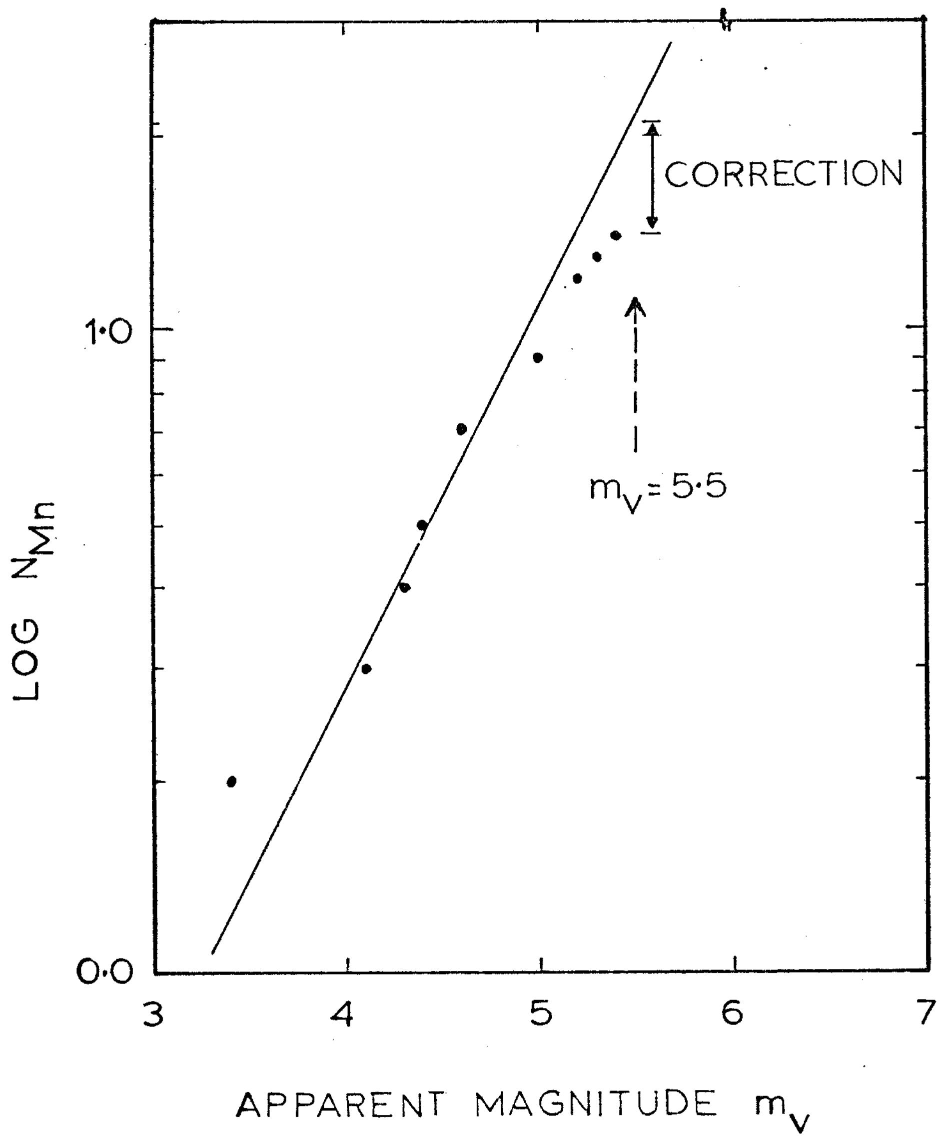


Figure II-2

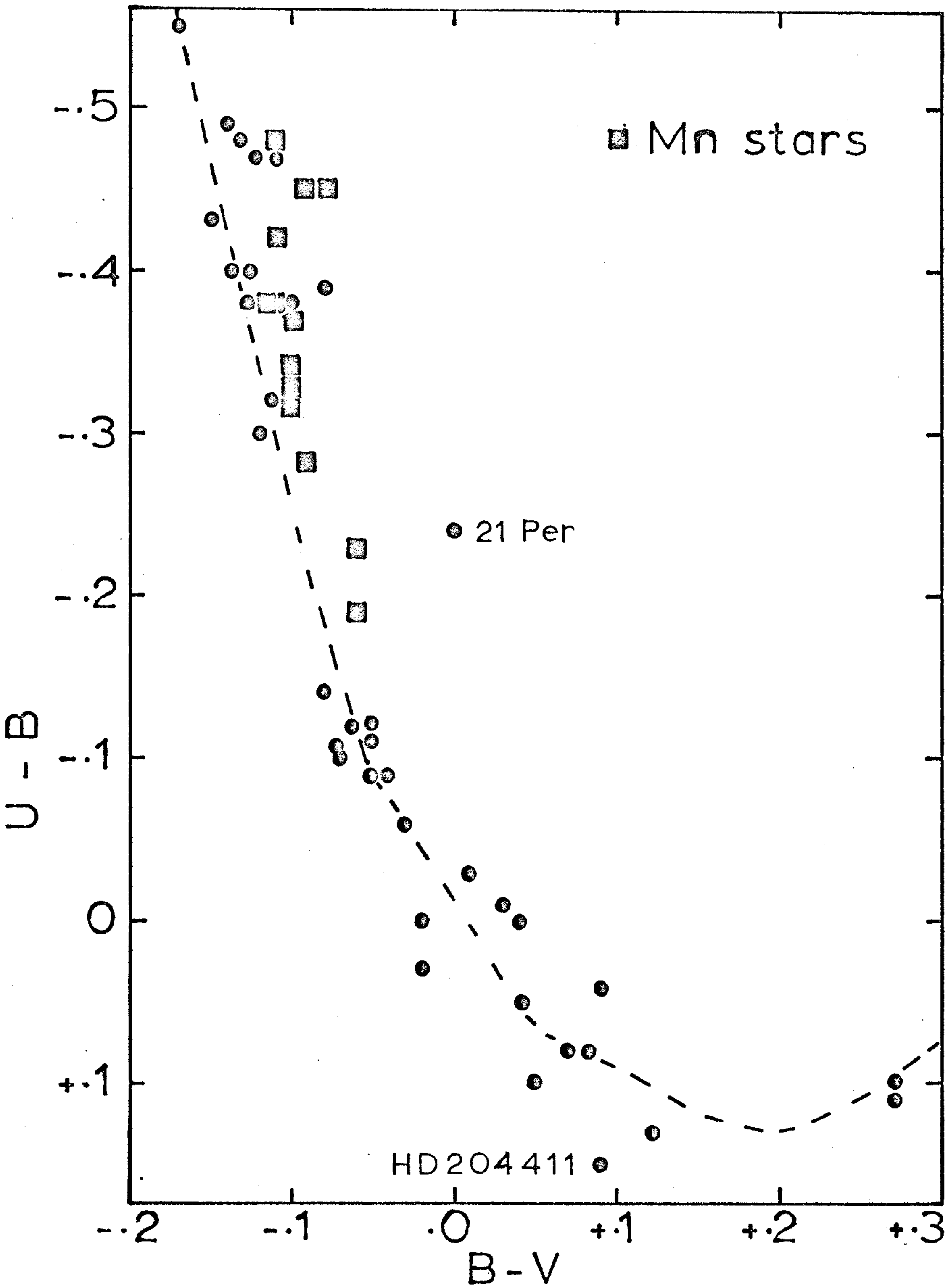


Figure II-4

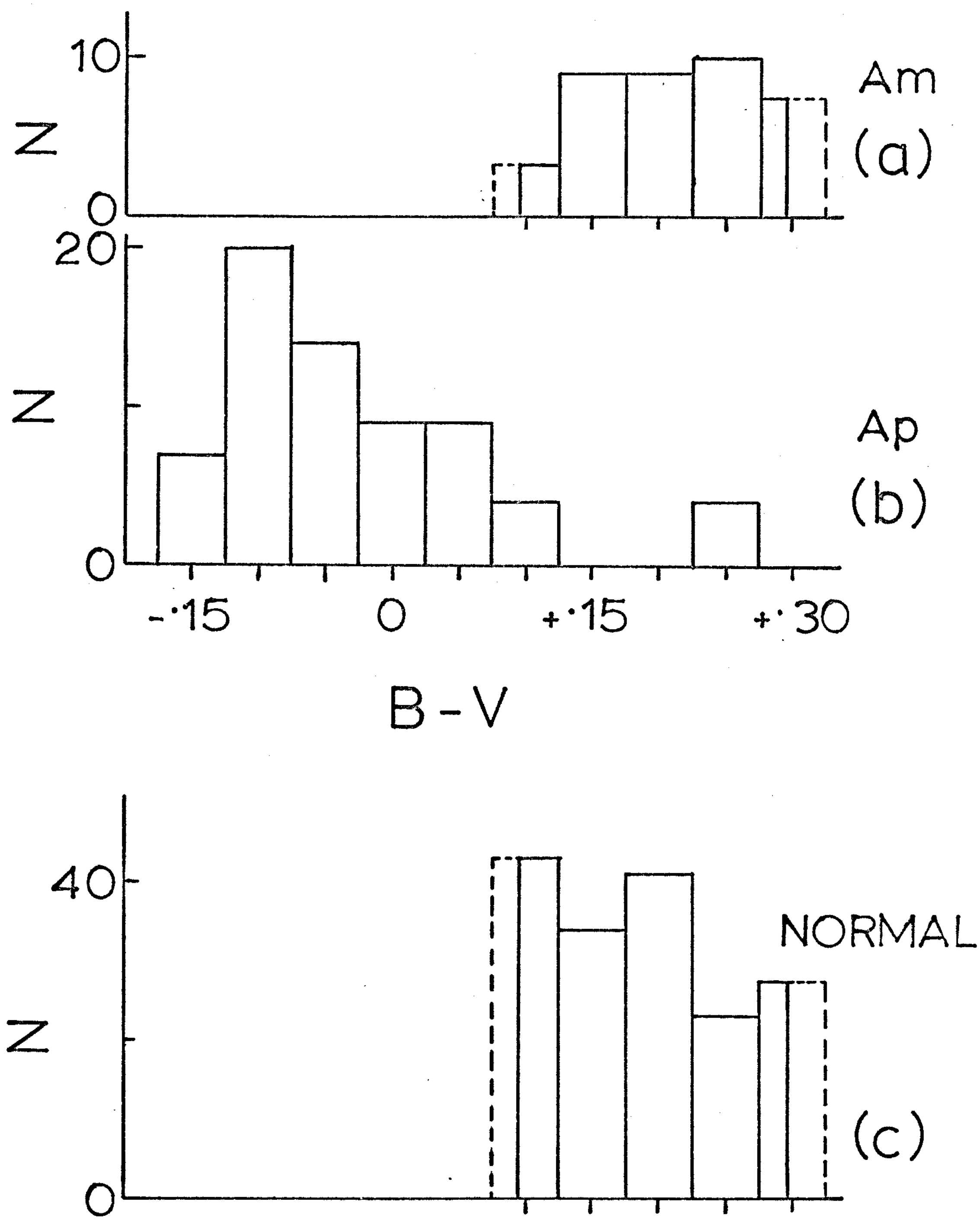


Figure II-5

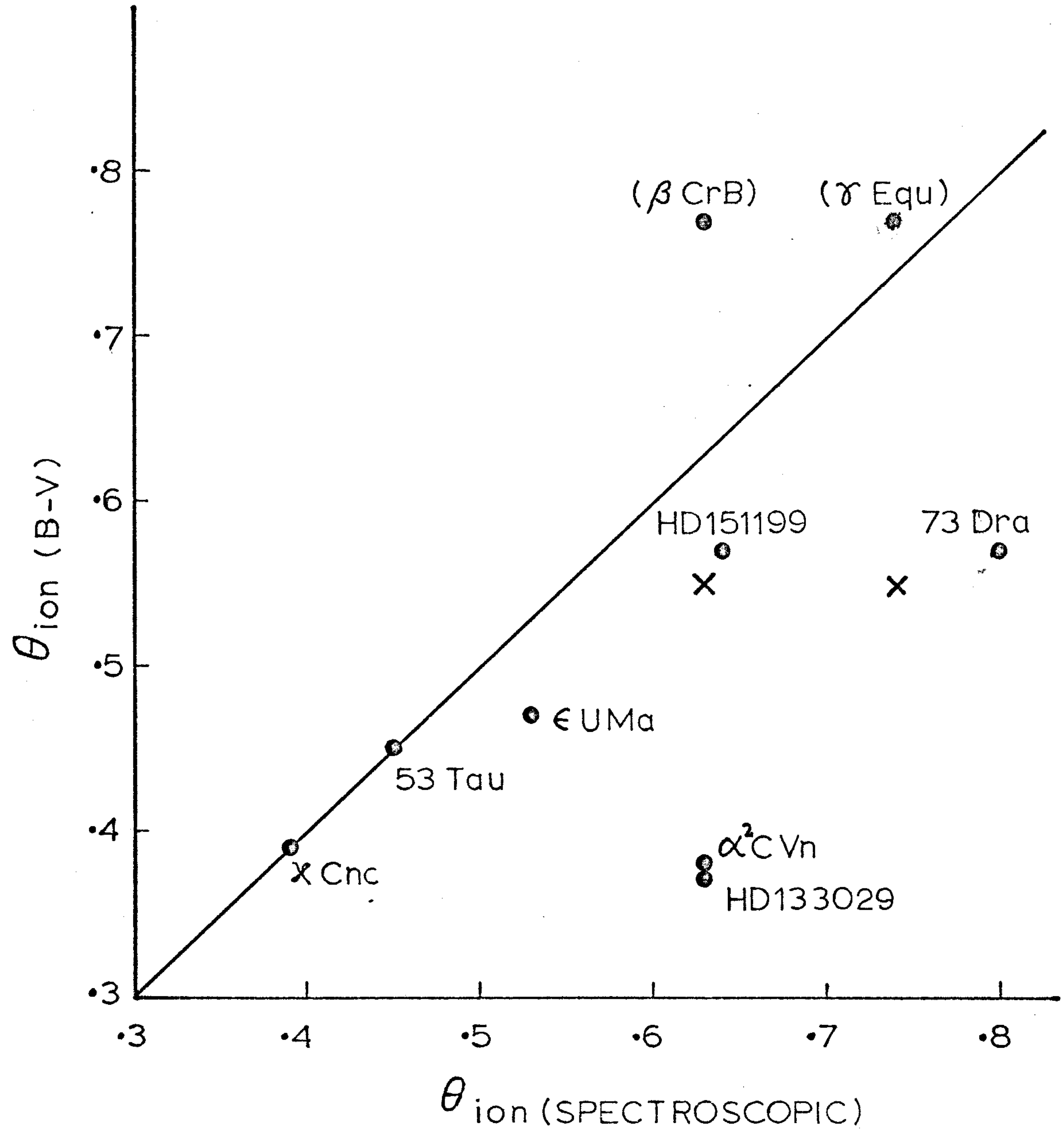

(β CrB)
(γ Equ)
HD151199
73 Dra
ε UMa
53 Tau
χ Cnc
α² C Vn
HD133029
•3
•4
•5
•6
•7
•8
θ ion (B-V)
θ ion (SPECTROSCOPIC)

Figure II-5

CHAPTER III

ULTRAVIOLET AND BLUE SPECTRUM

Section III-A

Problems of Ultraviolet Studies

While the "visual" and infrared regions receive less attention because of the slower speeds of the sensitized emulsions needed for these regions, study of the UV spectrum is handicapped by large atmospheric extinction of radiation. Atmospheric extinction increases very rapidly towards the shorter wavelengths, and as the wavelength 2950 Å is reached, the atmosphere becomes completely opaque. This causes difficulties in two ways: (a) much longer exposure times become necessary, and (b) rapid increase in atmospheric extinction produces a steep gradient in the energy distribution which permits only a rather small range in wavelength to be properly exposed on a single spectrogram. Matters are further aggravated for A-type stars because of a Balmer jump of more than one magnitude. Other factors which become important in practice are: (a) Atmospheric dispersion produces a small spectrum of the star's image on the slit of the spectrograph; for maximum efficiency, the invisible ultraviolet part should be centered on the slit. Consequently, "guiding" has to be done by, so to speak, an off-set method. This effect becomes important if the "seeing" is good. (b) The Sec Z factor (Z is the zenith distance of the star) has to be taken into account for estimating exposure time. (c) Losses of light in the ultraviolet make it imperative to use very fast cameras with short focal ratio. This, naturally, impairs the image quality compared with a camera having equal focal length but slower focal ratio.

It is for such reasons that a systematic study of the ultraviolet region of the spectrum of Ap stars has not received attention. However, our continued lack of an understanding of the nature of the Ap stars makes it desirable that we explore new areas to extend information.

Organization of Observations

As will be discussed in Chapter IV, a survey of the BeII doublet at λ3130-31 was a major motivation for undertaking the study of the ultraviolet spectrum. Use of the glass optics was, therefore, precluded. The 16-inch focal length camera of the 100-inch coudé spectrograph was used. This camera has a quartz field-flattener - the longer focal-length cameras do not have field-flatteners at all, but their use was not feasible due to a limitation of observing time, and the necessity of observing a large number of stars. Other guiding considerations for this program were:

(a) Since the stars of the present study are abnormal objects, it would not be safe to draw any generalizations from just a few stars. Therefore, it was decided to take spectrograms for <u>at least</u> half a dozen stars from each of the five currently known peculiarity groups, viz., "Mn", "Si-4200", "Eu-Cr", "Si-Cr-Sr", and "Sr".

(b) Earlier studies of Ap stars - Morgan (58), Searle and Sargent (69) - have dealt mainly with bright stars. In the present study, a large number of stars, fainter than $5^{m}_{.}5$, were included. The brighter stars of the present study provide a convenient overlap for comparing consistency of results with those of the earlier investigators.

(c) The ultraviolet spectrum of Ap stars is quite heavily crowded with spectral lines, and if a star has large rotational velocity (v sin i), excessive blending is produced and most of the spectral lines are "washed out". Therefore, most of the stars chosen have an estimated $v \sin i \lesssim 100$ km/sec.

(d) For selecting stars for observation, an extensive list was prepared. Then, subject to the above three criteria, stars close to meridian at the time of observation were picked.

Observational Material and Reductions

Observational material, on which this study is based, is described below. (A list of plates is given in Appendix II.)

(a) Thirty-four high dispersion plates (mostly 10 Å/mm or 4.5 Å/mm) of 21 Ap stars, and nine plates of five normal stars obtained from the plate file of the Mt. Wilson and Palomar Observatories.

(b) Plates taken expressly for this study: (i) Ninety plates taken with the 16-inch Schmidt camera of the Mt. Wilson 100-inch coudé spectrograph. For most of the stars, two plates were taken, one exposed for λ 3400 Å and the other for 3300Å $\lesssim \lambda \lesssim$ 3700Å. (ii) Ten plates taken with the 32-inch Schmidt camera of the same coudé spectrograph. Some of these plates were exposed for blue and others for UV. In both cases, (i) and (ii), grating 46B, having 15000 lines/inch and "blazed" for λ3500 in the second order, was used. The resulting dispersions in the UV are for (i), 21 Å/mm; and for (ii), 10 Å/mm. Eastman spectroscopic emulsion type IIa-O, was used. With this plate and grating combination, the entire range from 3000Å - 5000Å is free from any problems of overlapping orders.

Prior to exposure, plates were "baked" for 72 hours at 50°C. Exposed plates were developed - often several plates simultaneously - for 15 minutes in Kodak D-76 developer at about 68°F. Small departures in the temperatures of the developer, mainly due to effects of humidity, were compensated by multiplying the development time by a factor of $2^{\left(\frac{68-T}{10}\right)}$, where T is the temperature of the developer on the Fahrenheit scale.

The commonly used developer, Kodak D-19, produces less chemical-fog than D-76. But, the latter was chosen because earlier experimentation showed that, (a) D-76 increases the effective speed of the IIa-O emulsion, at low density levels, by at least 30%; (b) D-76 produces a finer grain; (c) it produces lower contrast (higher contrast is not desirable because it would further enhance the effect of the steep gradient of energy distribution which is already too steep due to atmospheric extinction); and (d) the long development time required with the D-76 makes control of effective development time more manageable.

On all the "16-inch" plates, the spectrum is 0.5 to 0.7 mm wide, and on the "32-inch" plates, 0.7 to 1.0 mm wide. The width of the spectrograph slit, projected on the plate, was 20 microns or less for all the plates. The star image was trailed, end to end, many times over the length of the slit so as to obtain the whole width of the spectrum uniformly exposed. Wavelength scale was calibrated on each plate by exposing iron-arc spectrum on both sides of the stellar spectrum. This calibration spectrum was exposed both at the beginning and at the end of the exposure of the stellar spectrum. In the path of the light beam from the iron-arc, an ultraviolet transmitting diffuser was used which enabled a calibration all the way down to 3100 Å or below.

A remark must be made regarding the image quality on these plates. The 32-inch camera produces excellent quality images over the entire length of the plate - the plate is curved in the plate holder, and it does not require a field-flattener. The 16-inch camera, on the other hand, uses a field-flattener and has a faster focal ratio. Furthermore, it does not have a provision for making a precise adjustment of the tilt of the plate holder for achieving a good focus over the entire length of the plate. The problem of focusing becomes further aggravated because this camera seems to be unstable and, sometimes, the position of the plate holder with respect to the focal plane changes capriciously. For these reasons, the quality of the 16-inch plates is variable; some are very good while others have a little "soft" focus. Nevertheless, all but a few spectrograms are usable.

Intensity Calibration

Intensity calibration in the UV presents difficulties that are not usually encountered for other spectral regions. In essence, the cause of the difficulty lies in the fact that all the ultraviolet absorbing elements in the light train cannot be dispensed with from the equipment available. Most important of all, hardly any radiation with wavelength shorter than $\lambda 3300$, is transmitted through the glass envelope of the photo-flood lamp. Due to glass optics in the wedge-spectrograph used for intensity calibration, the light intensity falls off rapidly below $\lambda 3800$ Å; in general, no reliable calibration curve could be obtained for $\lambda < 3700$ A (any inordinately long exposure time would only increase the scattered light). With the step-slits of the coudé spectrograph this limit can be pushed a little farther, to about $\lambda 3500$. But with the 16-inch camera, the strips impressed on the plate become too narrow and too closely spaced for derivation of a reliable calibration curve. The following solution was adopted ultimately:

On two different observing runs, calibration plates with step-slits, in conjunction with the 32-inch camera of the coudé spectrograph, were taken and calibration curves at $\lambda\lambda 3500$, 3700, and 3900 were determined. No systematic variation in these curves, as a function of wavelength, was found. On this basis, it was assumed that the calibration curve derived for $\lambda 3700$ can be used for the whole of the range of the ultraviolet spectrum ($\lambda 3700$ - $\lambda 3100$) under present study.

About three wedge-plates were taken for each night and were developed together with the plates of stellar spectrum. Exposure times for stellar spectra were typically around one hour and effects due to reciprocity-failure are minor (especially for the "baked" plates); however, the wedge-plates were suitably exposed to take this effect into account.

Microphotometer tracings, reduced to intensity scale, were made for all the plates using the Caltech Intensity Converter. Dispersion on the tracings was about 1/4 inch/Å - a magnification of about 130. The width of the scanning slit, projected on the plate, was 15 to 20 microns; the time-constant of the amplifier was less than the time taken by the plate-carriage to traverse 20 microns. Tracings, thus obtained, were used for the measurement of equivalent widths of selected lines and also for identification of lines. For the latter purpose, tracings were matched against a similar tracing for a standard star. Most of the plates were traced twice (As it turned out, on the earlier set of tracings, lines were erratically displaced, due to non-uniform motion of the plate-carriage of the microphotometer caused by dried-up lubricant.).

Section III-B

Identification Procedure

For UV Plates: The identification of lines in the spectrum of a star is a straightforward matter if the star has very sharp lines and accurate wavelengths of the lines are measured on high dispersion plates. Then, a great majority of lines being unblended, one can find the identification by using tables such as Moore's Revised Multiplet Tables (RMT) (56) or the MIT Wavelength Table (39). None of these fortunate circumstances was applicable in the present case, however, and vagaries of line strengths and plentitude of blendings on the present plates (21 Å/mm) made the problem several-fold complicated.

It took a circuitous path of trials and failures to evolve an expeditious procedure and experience for judgement for the identification of the lines and elements in the UV spectrum of a given star. There are two aspects of a program of line identification:

(a) Identification of all the noticeable features in the spectrum, and

(b) Development of criteria for estimating the weakening or strengthening of lines due to various elements for the study of which the UV spectrum is well suited.

For the Ap stars, the first of these, (a), becomes easier after the second, (b), has been established. But, for the establishment of criteria in (b) we need identification of lines in the vicinity of the criteria establishing lines, which are themselves not known a priori (it's a vicious circle).

While handling the type of spectrograms (and tracings) under discussion, it was necessary to be able to judge which details on the tracing were real and which were spurious (caused by grain, etc.). This judgement was developed by (a) comparing spectrogram with its tracing, and (b) by comparing tracings of different spectrograms of the same stars, or (c) by comparing tracings of the upper and the lower halves of the same spectrogram.

Finally, a very important consideration was that the identification procedure should be such that identification of lines on a very large number of plates, under this program, becomes feasible.

An identification procedure was ultimately evolved, for the 3700-3100 A region, as follows: On the basis of earlier trials it was found that, on the present set of spectrograms, lines due to Ca, Sc, Ti, V, Cr, Mn, Fe, Co, Ni, Y, and Zr had a chance of being noticeable in normal stars. (Of these, only the CaII lines were unsuitable due to serious blending in the Ap stars.) Elements Be, Mg, Si, P, and Eu were also included in this list because of the special interest in these for Ap star work. Then, on a tracing of a 10 Å/mm spectrogram of the sharp line A1V star o Peg, a wavelength list of about 10-15 strongest observed (or predicted) lines of each of these elements were marked (for Cr, Mn, Fe, and Ni, lines due to both neutral and singly ionized atoms were marked; for the other elements, only for the singly ionized atoms.). Laboratory intensity values and excitation potential of the lower level, as given in the RMT, were also noted. Next, to find actual or potential blendings, other weaker lines, due to the above listed elements, in the vicinity of the stronger marked lines, were marked on the tracing. Many spectral lines which are resolved on this spectrogram (10 Å/mm) become blended on a spectrum of the same star taken at 21 Å/mm. Therefore, in many cases, lines that are very suitable as criteria establishing "key lines," on 10 Å/mm plates, had to be rejected for work with 21 Å/mm plates. Information from the tracing of the 10 Å/mm plate was, then, transferred on a tracing of a 21 Å/mm plate of this star o Peg. This latter tracing of o Peg we shall call the "master" tracing. This "master" tracing was, next, further compared with similar tracings (i.e. with the same dispersion) of other Ap stars to check to see if any of these criteria establishing "key lines" were blended with lines of any of the elements greatly overabundant in the Ap stars. Lines selected after these screening procedures are listed in Table III-1. Some of these may still be blended, but these represent the best lines available in the UV, and suitable for work at this dispersion.

The strength of the lines in Table III-1 was noted, on a qualitative scale, for all the Ap and normal stars. From this survey 22 lines, which covered the different elements and represented most consistently the behavior of different elements, were selected for measurement of equivalent width. These 22 lines are indicated by an underscore in the column "Principal Contributor" in Table III-1. These 22 are, then, the "key lines", checking these on a tracing of a star readily tells us which lines of which elements are strengthened and of which elements are weakened relative to a normal star. Having this important information, a line-by-line identification for the rest of the spectrum of an Ap star can be made without major surprises.

Identification Procedure for High Dispersion Blue Plates

For the A-type stars, the blue region of the spectrum is not as crowded with lines as the UV region. Furthermore, the situation very much improves with the use of high-dispersion spectrograms. Consequently, the "key" or "test" lines could mostly be located by a visual inspection of the plates through a comparator or a magnifier. In a few cases of uncertainty, the wavelengths of lines on the plate were measured.

The procedure for the selection of the "key" or "test" lines was, in principle, the same as described for the UV spectrum, but not so difficult in practice. As a starting point, identifications of lines in the region $\lambda 3840$ - $\lambda 4640$ for o Peg by Maestre and Deutsch (48) were used. To this were added, on a high magnification photographic enlargement of a 10 A/mm plate of o Peg, "key" lines of other elements not observed in the spectrum of normal stars of type A.

Section III-C

Observability of the Various Elements

In this section, a brief account of the most suitable lines for detection of various elements, in the ultraviolet and blue regions of the spectrum, is given. By "UV" we shall mean the observable region on the short wavelength side of the Balmer discontinuity ($\simeq$ 3700 Å) - but mostly redward of $\lambda 3100$ or $\lambda 3150$ - and by "blue", the region on the long wavelength side of it up to $H\beta$. The first and second ionization potentials of the elements are designated by I and II respectively, and are in electron volts.

Be I = 9.3 II = 18.1

The most promising lines for the observability of Be are the λ3130.42 and λ3131.06 lines of the $2^2S - 2^2P°$ doublet of BeII. In normal A0 stars, the contribution to equivalent widths by these lines is of the order of 10mA. Further details on Be and other rare light metals are given in Chapter IV.

Mg I = 7.6 II = 15.0

There are no convenient lines of MgI or MgII in the UV. The close doublet MgII 4481.13 - .33 is the most convenient feature for a quick visual inspection. Throughout the range of Ap stars it is conspicuous, even at moderately low dispersions provided Mg is not underabundant by a large factor. For cooler Ap stars, MgI λλ3832.30, 3836.76, and 4167.27 are useful.

Si I = 8.1 II = 16.3 III = 33.3

For the hotter Ap stars, the best lines for detection of Si are the SiII lines λλ4128.05, -30.88 and λλ3853.66, -56.02, -62.59. Normally no Si lines are observed in the UV; however, in stars like HD224801, in which the usual SiII lines are very strong, SiII 3333.16Å is weakly present.

For normal stars, lines of SiIII do not appear in cooler than early B-types; however, in hotter Si-stars, like HD34452, 108 Aqr, τ^9Eri, HD224801, etc., SiIII 4552.56 is noticeable.

In the cooler Ap stars, lines of SiII become weakened and blended with other lines and, therefore, SiI lines should be consulted for more reliable estimates. But, of the two SiI lines in the blue, λ4102.93 lies close to H_δ, and λ3905.53 lies near CrII 3905.64, which in Cr-rich stars becomes dominant. Suitable lines of SiI lie in the infrared and "visual" regions.

<u>P</u> I = 10.9 II = 19.6

Observable lines of PI lie only far in the infrared; lines of PII are noticeable in the early B-type stars. Bidelman's (14) identification of PII lines in the B8 Mn star κ Cnc has made this element of interest for the Ap stars. The most important line of PII is λ4475.26 which, if present, can easily be located because of its proximity to MgII 4481.

<u>Ca</u> I = 6.1 II = 11.8 III = 51.0

The striking weakening of CaII lines, particularly of CaII-K, is the single most common peculiarity in the spectra of Ap stars. Except for the temperatures of the cooler Ap stars, the resonance line CaI 4226.73 is not expected to be detectable because practically all the calcium is ionized. Accessible lines of CaIII originate from 30eV levels and have not been identified, with certainty, in any of the normal B or A stars (Merrill (51)). However, since there is a possibility that CaII lines might be too weak due to excessive double ionization of Ca, the strongest CaIII lines, which are in the UV, were searched in some of the stars in which the CaII lines are extraordinarily weak. No strong lines of CaIII are present though the possibility of weaker lines cannot be ruled out because on the present plates blends could not be resolved.

Sc   I = 6.7   II = 12.8   III = 24.6

The ultimate lines of ScII, λ3613.84, lies between CrII 3613.2 and FeII 3614.87. If lines are not too wide and if CrII lines are not greatly strengthened, this line of ScII is quite satisfactory. Similar is the case with ScII 3630.74 and 3353.73. If the CrII lines are too strong, however, some estimate can be made from ScII 3572.52 (+ ZrII 3572.47 + NiI 3571.87), and ScII 3589.64, - 90.48 (+VII 3589.76).

In the blue, the strongest line of ScII is λ4246.82 while for visual inspection λλ4314.08, -20.75, and -25.01 are most easily identifiable.

In hotter Ap stars, lines of ScII become very weak, and nothing can be said about the over- or under-abundance of Sc, unless lines in that star are very sharp and high dispersion is used - e.g. χ Cnc. The two lines of ScIII, listed in the RMT, are not visible on a 10 A/mm plate of the sharp line star ϕHer (B-V = -.05), which has quite strong ScII lines for its color.

<u>Ti</u> I = 6.8 II = 13.6

There are no important lines of TiI in the UV or in the blue spectrum of the Ap stars. Lines of TiII occur copiously in the UV spectrum and, together with the lines of CrII and FeII, constitute a majority of the conspicuous features in the UV spectrum of the A-type stars. Lines TiII 3685 and TiII 3349 (plus a small contribution by CrII 3349.34) are easiest to identify for a quick visual inspection. Other useful lines for making visual estimates are: $\lambda\lambda$3383.76 and 3329.5 in the UV, and $\lambda\lambda$4468.49, 4171.90, 4163.64, etc. in the blue.

<u>V</u> I = 6.7 II = 14.1

Lines of VII are usually weak both in the UV and in the blue, and, like ScII lines, become difficult to observe in the hotter Ap stars. There are no very good lines in the UV for detection of VII with 20 A/mm. The strongest lines of VII are $\lambda\lambda$3093.11 and 3102.30 which are not easily accessible; the former is seriously blended with CrII. Other useful lines are $\lambda\lambda$3589.75, 3545.19, and 3271.12.

In the blue, VII 4005.71, -23.29, and -35.63 are the convenient and blend-free lines, though not among the strongest.

Cr I = 6.7 II = 16.6

In the UV, lines of CrII are even more numerous than those of TiII, and in those Ap stars which have chromium abnormally strong lines of CrII alone constitute the majority of the noticeable lines.

The CrII line λ3677 which is most easily reconizable in all the A-type stars, and the other strong CrII lines, λλ3632, 3603, and 3585, form a pattern which is outstanding in the Cr-stars (Figure III-1). A few lines of CrI are present only in the late A-type stars; in Cr-stars, CrI 3578.68, 3593.49, and 3605.33 become quite strong - these are blended, however.

In the blue, useful lines are CrII λλ4558.66, 4261.92, CrI4254.36, etc. Although the blue region has many chromium lines, the profusion of strong lines is not so great as in the UV. One must be particularly careful of blends with chromium lines in Cr-stars.

Mn I = 7.4 II = 15.6

The strongest lines of MnII are those of the multiplet $a^5D - z^5P^o$, in the UV. These are present with fair strength in the normal stars from late B to middle A-types. In the Mn-stars, which are usually hotter Ap stars and do not have crowded spectra, these lines are very prominent, and even at moderately low dispersions can be detected easily. In the blue, on the other hand, weak MnII lines are seen in the Mn-stars only. Of these, $\lambda\lambda$4136.96, 4206.38, 4292.25, and 4326.76 can be seen in most of the Mn-stars. On high dispersion spectrograms of the very sharp line star χ Cnc, Bidelman (14) finds "of the 244 lines measured, 110 can be attributed, wholly or in part, to MnII."

Resonance lines of MnI, λ4030-34, begin to appear in normal stars cooler than A0. These lines are also noticeable in the cooler Mn stars ϕHer and ι CrB (B-V = -.05 for both.) but not in the hotter members of the Mn-group. In Cr-stars, blending of these MnI lines with chromium lines can render estimates erroneous, e.g. in HD153882 the apparent strengthening of MnI lines (5) is due to Cr lines.

<u>Fe</u> I = 7.9 II = 16.2 III = 30.5

In the UV there are fewer strong FeI lines than in the blue, though the ultimate line λ3581.20 and another good line λ3570.10 lie in the UV. The FeI line 3565.38 lies close to FeII 3566.15 and, if the lines are sharp, these are useful in estimating FeI/FeII ratio. For a similar visual estimate in the blue, convenient pairs are FeI4415.41, FeII4416.82; and FeI4383.55, FeII4385.38.

Lines of FeII are numerous both in the UV and in the blue. The region between λ3150 - λ3300 is especially rich in strong FeII lines, which in the spectrum of η Leo (AOIb) become dominant in that region.

Lines of FeIII, in the UV or in the blue, are not seen even in the hotter of the usual Ap stars.

Co   I = 7.8   II = 17.1

For observing cobalt, the UV region is more suitable than the blue. According to the RMT, the strongest of the CoI lines lie in the UV, and both of the two multiplets of CoII listed have lines in the UV. However, all these lines are weak and blended (at 21 A/mm). The only line from which reasonably reliable information on the presence of cobalt can be derived is CoII 3501.73 (in HD9996 this line is blended with a strong unidentified line at λ3502.1, which might be due to some rare-earth element).

Ni   I = 7.6   II = 18.4

The UV region has many more of the stronger NiI lines than the blue and these lines are easily seen in a sharp line star like υ Peg. The strongest line of NiII, λ3087.02, is free from blends only at high dispersion and if the lines are sharp. The next strongest line of NiII, λ3513.93, begins to be contaminated with FeI 3513.82 at AO.

In the blue, NiII 4067.05 is the most convenient line.

Ga

Bidelman and Corliss (15) have identified the following four lines of GaII in the spectrum of χ Cnc.

Wavelength (Å)	Intensity in χ Cnc
4251.15	1 - 2
4254.09	0 - 1
4255.71	2 - 3
4262.00	3

In practice, λ4255.71 is found to be most suitable for detection. The CrII line 4261.92 is present even in the normal star θ Peg and becomes very strong if the star has an overabundance of Cr. Therefore, unless accurate measurements are made, GaII 4262:00 is not safe for detection.

Sr I = 5.7 II = 11.0

The zero volt lines of SrII 4077.71 and 4215.52 have the distinction of being strengthened and observable, to a varying degree, in Ap stars of all the types. In normal spectra of types F to M, the intensity of these lines has a strong positive correlation with the luminosity of the star. But, in A-types, and particularly in the early A-types, population of singly ionized strontium atoms has passed its maximum, and a decrease in electron pressure will reduce the fraction of Sr atoms in the singly ionized state. Consequently, in the Ap stars, strengthening of SrII lines cannot be attributed to the luminosity effect.

Merrill and Lowen (52) have discussed blending with the SrII lines in the late-type stars. For Ap stars, however, the possible blendings are totally different, and are given below:

SrII	4215.56	(300)	SrII	4077.71	(400)
CrII	15.77	(2)	SiII	76.78	(1)
CrII	17.07	(1)	SiII	75.45	(2)
LaII	17.56	(200)	CrII	77.50	(4)
			CrII	76.87	(3)
			LaII	77.35	(300)
			DyII	77.98	

The numbers in parentheses are intensity values as listed in the RMT. At high dispersion most of these lines are resolved, but at low dispersions these lines blend with the SrII lines, which form classification criteria for low dispersion work. The SiII lines make an important contribution in the Si-4200 stars. In stars with CrII lines outstanding, the CrII contribution to λ4077 becomes considerable, but the contribution to λ4215 is less important. The rare-earth lines will make a significant contribution only if these rare-earth elements are unusually overabundant. SrII 4215 is, therefore, more reliable for identification at low dispersion.

In the Sr-stars the UV lines of SrII, λλ3380.71, 3464.46, and 3474.89, are present but blended with other lines.

<u>Y</u> I = 6.5 II = 12.3

In normal stars YII lines begin to appear at early A-types (Merrill (51)). But in the Mn-stars (usually about B8 and B-V = -.10) many YII lines are seen both in the UV and in the blue. The ultimate line of YII lies on the wing of Balmer H15. Convenient and relatively blend-free lines in UV are λλ3600.74, 3633.12, and in the blue λλ3774.33, 4374.94 (plus ScII 4374.46), 3950.35, 3982.59, etc.

<u>Zr</u> I = 6.92 II = 13.97

Lines of ZrII are quite weak in the early A-types, but strengthen in the later types. The UV region has more of the stronger ZrII lines than the blue, but at 21 A/mm most of these are blended. The ultimate line λ3391.96 is usable only at high dispersion, consequently, λ3438.98 is usually the most reliable line. ZrII λλ3698.17 and 3709.27 fall on the wings of hydrogen lines H17 and H15 respectively.

Convenient lines in the blue are λλ4149.22 (I = 75), and 3958.24 (I = 50). But in the rare-earth rich stars these may be blended with NdII and TmII respectively. In that case, weaker lines λ3991.14 (I = 40) and λ3998.98 (I = 30) should be used.

Ba   I = 5.19   II = 9.96

In Ap stars, Ba is either weak or normal. Therefore, only the resonance line BaII 4554.03 is useful for detection of Ba. The other line BaII 4166.00 becomes noticeable only in the cooler normal A-type stars.

Rare Earths

The presence of lines due to the rare-earth elements in the spectrum of many Ap stars is well known. Of these, lines due to EuII are most conspicuous. In HD188041 (B-V = +.20; A5p) at EuII-maximum phase, lines of EuII are extraordinarily strong and are next to only the Balmer lines and the CaII K & H! No important lines of EuII lie in the UV, however. Lines of CeIII and DyII are quite strong in the UV in some stars, but at 21 A/mm, dispersion of the plates used for this survey, most of these are blended. For a systematic study of the rare-earths, therefore, some plates were taken with 10 A/mm, both in the UV and in the blue, and results on all the rare-earth elements are given in Chapter V.

All the UV lines used for identification work are listed in Table III-1. In this table, relevant neighboring lines, which might blend with the principal line if the dispersion of the plate is not high, are also given. Lines not noticeable in the spectrum of the sharp line AIV star o Peg but which may become significant in some Ap star, are indicated by (*) and other weak lines with equivalent width about 10-20 percent of that of the principal line are indicated by (#). The blue lines, used with the high dispersion plates, are listed in Table III-2.

Section III-D

"Abundances"

The question of whether the observed spectral peculiarities are due to some unusual physical conditions in the atmosphere of the Ap-stars, or to real abundance differences has been investigated by various investigators over the past 25 years. However, some of the results of even the recent investigations are in sharp disagreement.

For a reliable determination of the abundances in Ap stars, some of the assumptions applicable to normal stars must be avoided. This imposes stringent requirements on the quality of observational material, and this requires one to be particularly careful in making analyses. For example, if enough reliable and weak lines (which lie on the linear part of the curve of growth) are not used, one cannot tell whether the strengthening of lines is due to an overabundance or due to an increase in micro-turbulence.

In the present study our aim is to find patterns of abundance anomalies among the various types of Ap stars - in all, for about 10 - 20 elements in 40 Ap stars. A detailed analysis for all these stars was neither necessary nor feasible. (It was for these reasons that a dispersion of 21 A/mm was considered acceptable; otherwise, to obtain meaningful results by the curve of growth method, a dispersion of at least 10 Å/mm is necessary.) However, by a careful selection of those lines from the spectrum which are insensitive to the uncertainties in physical conditions, we can learn about the intrinsic behavior of that element.

These behaviors and patterns of correlations can be studied at levels ranging between these extremes: one, in which no allowance is made for the differences in the physical conditions among various stars and correlations are derived from the observed line strength (equivalent width) only; and the other, in which correlations are derived from truly intrinsic abundance anomalies. The atmospheres of the normal stars are presumed to be describable by specifying the effective temperature, gravity, and chemical composition. By accounting for the effects of differences in (a) effective temperature, (b) electron pressure, (c) continuous absorption coefficient, and (d) micro-turbulence, one can determine the intrinsic abundance differences between two stars. In the case of Ap stars, however, there may be additional phenomena which alter the strength of the spectral lines. Consequently, after accounting for the effects of the four parameters listed above, the abundances determined may still not be intrinsic.

In the present chapter, for determining whether a certain element is overabundant or underabundant, we shall examine only the parameters mentioned above. (In view of the remarks made in the preceding paragraph, to reduce the possibility of misinterpretation, in the subsequent discussion we shall refer to a given element as "strong" or "weak" instead of the conventional terms "overabundant" and "underabundant".) These estimates have been made on a qualitative scale as will be described later. It is to be noted that the fact that the results concerning "weakening" or "strengthening" derived in a manner described in this chapter might not indicate "intrinsic" abundance anomalies in some cases, does not decrease the usefulness of these results. For example, as will be discussed later, the cause of the strong positive correlation between the "weakenings" of Ca and Ba (derived from CaII and BaII lines) might not be an intrinsic underabundance of Ca and Ba. If such is the case, the value of this correlation lies in pointing out the existence of a phenomenon not yet discovered.

Now we discuss the manner in which the effects of the earlier mentioned parameters have to be taken into account or eliminated for making estimates.

A. Temperature

In Section II- B we have discussed the (B-V)-temperature relation for the Ap stars. To take the temperature differences into account, therefore, all the stars, Ap and normal, were arranged in a B-V sequence (Appendix II) and the spectrum tracing (or plate) of a given Ap star was compared with the tracing (or plate) of normal stars available on the two sides of the B-V value of the Ap star. The working of the procedure can be seen from the various illustrations reproduced here in which spectra of the various stars (normal and peculiar) are arranged in a B-V sequence. (See e.g., Figures VI-1 to 5.)

It is to be emphasized that our aim is to find which elements are too "strong" and which are too "weak", rather than to find their actual abundances. Therefore, instead of assigning a precise value of 0 to each star, estimates were made by adopting such a comparison that a change in in the direction of the more appropriate value will only enhance the degree of abnormality. For example, for the Ap stars HD34452 and 108 Aqr, spectra of the normal stars Her and πCet (B-V = -.17, -.15, -.18, and -.14 respectively) were available for comparison. In Figure VI-2 we notice that in these Ap stars FeII4122.64 and 4233.17 are much stronger than in πCet, in which

FeII4122.64 is only barely visible. Similarly, CrII4242.38 is very evident in the Ap stars but absent in πCet. Therefore, even if we use a comparison for which B-V color is cooler than that for the Ap stars, we can easily conclude a "strengthening" (or overabundance) of Fe and Cr in these Ap stars. If a comparison with appropriate B-V is used, the degree of abnormality will only be enhanced, because at the temperatures of these stars population of the singly ionized atoms has passed its maximum, and any further increase in temperature will decrease the fraction of the atoms in the singly ionized state.

B. Electron Pressure

The electron pressure value for each star, Ap and normal, was derived using the Inglis-Teller formula:

$$\log P_e O = 11.11 - 7.5 \log N_m$$

Values of N_m, the upper quantum number of the "last visible" Balmer line, were derived by drawing upper and lower envelopes to the Balmer series near confluence point, on the spectrum tracing, and noting the convergence point. Values of mean N_m and $\Delta \log P_e$, the departure of electron pressure value relative to a normal star of the same B-V value, are listed in Table III-3. Only four

Ap stars of this survey, HD192913 , OHer , HD204411 , and ρ Her(A) show significant departures with $\Delta \log P_e = -.20$, $-.20$, $-.35$, and $-.30$ respectively. Further examination of their spectra showed that only HD192913 and ϕHer are genuinely peculiar stars. The other two have only a low degree of peculiarities which, probably, can be explained by taking the low P_e value into account. Consequently, results for the latter two stars are not presented in Table III-4. For HD192913 and OHer effects of low P_e are small but noticeable. Effects of these departures were taken into account by referring to the spectrum of A0Ib star η Leo and estimating the effect of decreased electron pressure on the various lines.

The electron pressure can effect the line strengths in another way: An <u>increase</u> in P_e can cause suppression of the double ionization of metals increasing the fraction of atoms in the neutral and the singly ionized states, thereby increasing the strength of the lines produced by the neutral and singly ionized atoms.

The effect of such a suppression of the double ionization is calculated for HD8441, in which copious lines of CrI and CrII are present; these lines are absent in normal stars with the same B-V value. For HD8441, B-V = 0 (therefore $\theta_e = .50$ and $\theta_{UV} = .61$; θ_{UV} is defined in Chapter IV) and $n_m = 19$ (these parameters do not indicate a higher-than-normal value of P_e). If we assume that the differential error $\Delta n_m = +1$ (this is certainly an upper limit to the error in n_m for work relative to normal stars and using 21 A/mm dispersion), the increase in the population of Cr^+ ions will be 2 percent for $\theta = .61$ and 24 percent for $\theta = .50$. With an increase in temperature, the absolute population of Cr^+ itself starts decreasing rapidly. Thus we find that even with the conservative estimates of the errors, the effect is not significant.

C. Continuous Absorption Coefficient

Neutral hydrogen is the dominant source of the continuous absorption for the range of Ap stars under consideration. Consequently, the 'abundances' inferred from the observed strength of lines are relative to hydrogen.

Because of the differences in exposure time, spectrogram quality, etc., often results are more accurate if instead of making a direct comparison of line strengths between two stars, intensities are measured relative to suitably chosen standard lines; for example, instead of comparing CrII4262 directly, the ratio CrII4262:FeII4233 in two stars is compared. In doing so, we compare abundance Cr relative to Fe in the two stars.

However, for the Ap stars, it could not be assumed *a priori* that a given element (Fe, say) has per unit mass the same abundance in all the stars. In fact, our survey clearly indicates the existence of significantly different Fe/H ratios in several stars. On the other hand, the method described in the beginning is free from such effects. For this reason, the first method was used as primary; though, for cross-checks, the ratio of line strengths were also estimated.

D. Micro-Turbulence

Other factors remaining the same, the effects of micro-turbulence on spectral lines falling on the different parts of the curve of growth are these: (i) Equivalent widths of lines on the linear part remain uneffected by any change in micro-turbulence velocity. (ii) Equivalent widths of lines on the flat part of the curve of growth increase (relative to a star with zero micro-turbulence velocity) by a factor $\simeq \left[1 + \frac{V_{tu}^2}{V_{th}^2}\right]^{1/2}$ where V_{tu} and V_{th} are the usual turbulence and thermal velocities.

(iii) For lines on the damping part of the curve of growth, collisional damping is the main source of broadening of the intrinsic profile and therefore the effect of microturbulence is less important.

These characteristics help us in estimating strengthening or weakening of the various elements in the following manner:

(a) If in an Ap star those lines of a given element which are absent or very weak (on the linear part) in a normal star, are present or strong, we can readily conclude an overabundance of that element. For example, in HD8441, the five lines on the short wavelength side of FeII4541.5, indicated by a bracket in Figure VI-1 , are all due to CrII or CrI. These lines, which have very low laboratory-intensity values, are absent in the sharp line normal stars. Hence, an overabundance of Cr in HD8441, by a large factor, is easily concluded.

(b) Strong lines such as MgII4481, CaIIK, and, in the genuine Si-stars, SiII4131 fall on or close to the damping part and, therefore, are not very sensitive to micro-turbulence difference. [Of these, though, λ4481 and λ4131 are high excitation lines; these are rather insensitive to physical conditions (temperature and electron pressure) over the range of the Ap stars under consideration for the reasons discussed by Searle and

Sargent(69).]

(c) If in an Ap star, lines due to a given element are weaker - on any part of the curve of growth - than in the normal stars which have zero or very small micro-turbulence, underabundance of that element in the Ap star can naturally be concluded. Weakening of FeII lines in some Mn-stars is illustrated in Figure VI-4 .

(d) The micro-turbulence velocity in the metallic-line stars is about 4 km/sec (quoted by Miss Underhill (76)). Also, the reliable analyses of Ap stars (Deutsch (29)) do not give turbulent velocities greater than about 4 km/sec. Furthermore, Kahn (46) has pointed out that micro-turbulence velocity cannot be greater than about one-third the velocity of sound, which is about 8-10 km/sec for the temperature range of the Ap stars. Therefore, if we take 4 km/sec as the limit for the turbulence velocity, we get

$$\left[1 + \frac{V_{tu}^2}{V_{th}^2} \right]^{1/2} \lesssim 1.75$$

Consequently, in an Ap star, if the equivalent width of lines, falling on the flat part of the curve of growth for normal stars, is increased by a factor greater than ~1.75, then that element is almost certainly overabundant.

(e) Some qualitative idea about the magnitude of micro-turbulence can be obtained by an inspection of the spectrum. If, for most of the elements, lines on the flat part are strengthened, it is quite likely that micro-turbulence will account, at least in part, for the strengthening of lines. This is the case with the metallic line stars. Also in the spectrum of Ap stars such as HD2453, or 17 Com A, such manifestations of micro-turbulence are obvious. On the other hand, if the intermediate strong lines of only some of the elements are strong, while those of others are normal or weaker then in the normal stars, the abnormal line intensities are probably due to causes other than micro-turbulence.

It is to be noted that the effect of magnetic intensification on the strengthening of lines is analogous to that of an increase in micro-turbulence: By magnetic intensification only those lines will be strengthened which lie on the flat part of the curve of growth.

After taking the above mentioned characteristics into account, the question of abundance estimates reduces to that of intensity estimates. Some relevant details concerning this process are: (a) In very sharp-line stars, quite weak lines can be observed, but in wide-line stars, all such lines are washed out. This effect was compensated by using a low-power magnifier for examining spectrograms of wide-line stars. Even for sharp-line stars, a power much more than X10 is not desirable. (b) A scale of 4 (sometimes suffixed by a + or - , and with 0 denoting normal, +4 very strong, and -4 very weak) was originally used for the qualitative estimates of "over- and underabundances." However, for convenience of tabulation, these values are presented on a scale of 12.

"Abundance" estimates have been presented in Tables III-4 (a-d). In the light of the present study, classification of some of the stars has been changed from that given in their original sources. (See Chapter VI)

Accuracy of the Results

At 21 A/mm, lines with equivalent widths less than 30-40 mA are subject to large errors. Experience has shown that, in such cases, visual comparison, on an intensity tracing, of a large number of such weaker lines gives more reliable estimates than calculations based on measurements of one or two lines for each element.

The high dispersion spectroscopic material for the blue region of the spectrum provided (in addition to information on elements which do not have suitable lines in the UV) independent checks for the estimates made earlier with the UV spectrum. From experience it was found that if the plates are not greatly over- or under-exposed, a visual examination of the high dispersion plates provides more reliable estimates than those made with intensity tracings of the lower dispersion plates. (If photographic calibrations for these high dispersion plates, obtained from the Mt. Wilson plate file, were available, more quantitative results would be possible.)

It is important to have an idea about the reliability of results obtained by this procedure. Accuracy of the results can be judged from a comparison of results obtained here with the published ones. For two Ap stars, earlier than A5, of this program - α^2 CVn and χ Cnc - published results are available. The present estimates and the published results are listed in Table III-5 and plotted in Figure III-2. Remembering that our aim is to find which elements are strengthened and which are weakened and to what degree (qualitatively), we find from this figure that the agreement is entirely satisfactory: All the points lie in the first and third quadrants in Figure III-4.

It was possible to make one more estimate of the reliability of this procedure by comparing the spectrograms of α CMa and α Lyr ; abundance analyses for these stars are available (Boyarchuk (19), and Hunger (41)). For different elements, i, estimates for abundance differences between α CMa and α Lyr relative to hydrogen, i.e. values of

$$\left[\log\left(\frac{\epsilon_i}{\epsilon_H}\right)_{\alpha\, CMa} - \log\left(\frac{\epsilon_i}{\epsilon_H}\right)_{\alpha\, Lyr} \right]$$

have been compared with the published results in Table III-6 and in Figure III-3 . Agreement is, again, very satisfactory. It is to be noted that the published results are absolute abundance determinations, and, have large probable errors ($\Delta \log \epsilon_i$ between .3 and .5). It would not be surprising, therefore, if it might account for some of the scatter, especially for Si and Al.

TABLE III-1

The Test Lines in the Ultraviolet Region of the Spectrum

λ	Principal Contributor	Intensity RMT	Neighboring Lines
3600			
85.19	TiII	250	
77.93	CrII	30	EuII* 78.26
.86		50	
.69		40	
74.74	ZrII	40	EuII*74.63
59.77	TiII	60	CeII*59.23
47.84	FeI:	100	CrII* 47.40
42.79	ScII:	40	CrII* 43.22
33.13	YII	200	CrII 34.04; CrII 33.45; FeII 32.29
30.74	ScII	50	FeI 31.46; (+CrII)
14.87	FeII	5	ZrII* 14.79
13.84	ScII	60	CrII* 14.26; CrII 13.2
11.06	YII	200	EuII 11.57; NiI 10.46
08.86	FeI	100	SmII* 09.49; NiI* 09.31; CrII* 08.66
05.33	CrI	1600	
01.93	YII	100	NiI* 02.28
00.74	YII	300	
3500			
97.71	NiI	50	
93.49	CrI	2100	
93.32	VII	600	
92.01	VII	800	GdII* 92.71
90.48	ScII	20	
90.46	SiIII*	8	
89.75	VII	1000	ScII# 89.64
86.54	MnI*	30	FeI 86.99; AlII* 86.56
84.53	YII	100	CrII 3585 + FeI 3585; GdII* 84.96
81.20	FeI	250R	ScII# 80.93
78.69	CrI	2400	ZrII# 78.22; FeI# 78.38
76.76	NiII	3	CeII* 77.46; ZrII# 76.88
76.34	ScII	35	DyII* 76.25
72.52	ScII	50	ZrII 72.47; NiI 71.87
70.10	FeI	100R	MnI# 3569
45.19	VII	1000	GdII* 45.80
32.12	MnI*	50	DyII* 3531.70
32.00		50	
31.85		30	

TABLE III-1, Cont.

λ	Principal Contributor	Intensity RMT	Neighboring Lines
3500, Cont.			
30.77	VII	500	EuII* 31.15; SmII* 30.60; (+ZrII 30.65?)
24.54	NiI	200	VII# 24.71; GdII* 24.20
15.05	NiI	150	FeII# 15.82
13.93	NiII	8	NiI* 13.93; FeI# 13.82
10.84	TiII	60	CrII 11.84; SmII* 11.23; NiI 10.34
04.89	TiII	80	Zr# 05.67, 05.47; GdII 05.51; VII# 04.43
02.28	CoI*	100	NiI 02.60
01.73	CoII	200	BaI* 01.11; NiI* 00.85
3400			
97.54	MnII	25	ErII 97.90; FeI# 97.84
96.81	MnII	20	VII 97.03
95.18	ZrII	50	YII 96.08
95.83	MnII	40	FeII 95.62; CrII 95.56; CrII 95.37
88.68	MnII	40	CrII* 89.45; CrII* 89.07; FeII# 87.99
74.12	MnII	50	CrII# 75.13; SrII* 74.89;
74.04		40	HoII* 74.26; FeII* 73.83
71.35	NiII	2	CrII# 72.07; CeIII* 70.89
68.68	FeII	8	GdII* 68.99; GdII* 67.26
			AlI# 58.23; CrII 57.62
54.16	NiII	5	CrII# 54.98; GdII* 54.90; CeIII* 54.37; GdII* 54.15
53.51	CoI*	200	TmII 53.67; FeII* 53.60; Ni* 52.89
46.26	NiI	100	CoII# 46.40; FeI# 45.15
38.98	MnII	20	GdII* 39.99, 39.78, 39.21; CrII* 38.46
38.23	ZrII	100	CrII* 37.93; NiI# 37.28
30.53	ZrII	30	TmII* 31.20; CrII* 30.42
14.77	NiI	150	ZrII* 14.65; FeII 14.14; NiI 13.94
12.63	CoI*	80	NiI# 13.48; FeI# 13.14
12.34		80	
07.30	NiII	8	GdII* 07.61; FeI# 07.46
05.12	CoI*	150	CeII* 05.98; ZrII# 04.84
03.32	CrII	100	
01.76	NiII	2	CrII 02.43; GdII* 02.07

TABLE III-1, Cont.

λ	Principal Contributor	Intensity RMT	Neighboring Lines
3300			
91.96	ZrII	100	GeI# 92.65; GdII* 92.53; EuII* 91.99; CrII 91.43
83.76	TiII	125	SmII* 84.66
82.68	CrII	60	SmII* 82.40
73.98	NiII	4	ZrII 74.71, TiII 74.35; NiI 74.22; CeII 73.73
70.94	CoII	50	SmII* 71.21; FeI 70.79
58.50	CrII	75	GdII* 58.62; FeII# 58.25
57.40	CrII	40	ZrII# 57.26; CrII* 57.22
53.73	ScII	25	CeIII* 53.26; CrII 53.12
50.42	NiII	5	GdII* 50.47
33.16	SiII*	2	
29.46	TiII	70	CrII* 29.46
27.89	YII	100	CrII 28.35
03.47	FeII	4	NaI 02.94, 02.34; EuII* 01.95
02.86	FeII	4	
3200			
89.35	FeII	7	YbII* 89.36
88.81	FeIII*	15	
77.35	FeII	9	EuII* 77.78
76.12	VII	1500	FeII 76.61; CrII* 75.92
76.08	FeIII*	15	
74.94	NiII:	3	CeII* 74.86
71.12	VII	1200R	TiII 71.65; CrII 70.14
67.71	VII	1000R	CrII 68.48; FeII 67.04;
66.88	FeIII*	20	FeII 66.94; EuII* 66.39
59.05	FeII	10	CrII 58.77
58.77	FeII	10	
55.88	FeII	8	CrII* 55.62; CrII# 55.30
30.55	SiIII*	3	SmII* 30.56; FeII 30.50
27.73	FeII	13	FeI# 27.80; CrII* 27.48; CeII* 27.11
13.31	FeII	13	CrII* 13.46; CrII# 12.91
10.45	FeII	10	SiIII* 10.52; SiII# 10.04
03.89	SiII	2	CrII 05.11; FeII 03.74; CrII 03.53
31.06	BeII		See Table IV-1
30.42			

* Lines not noticeable in the spectrum of the sharp line A1V star ο Peg, but may become significant in Ap stars in which lines of that element are particularly strong.

\# Weak lines with equivalent width 10-20 percent (or less) of that of the principal line in ο Peg.

TABLE III-2

TEST LINES IN THE BLUE REGION OF SPECTRUM

3774.33	YII	4246.83	ScII
3853.66	SiII	4247.43	FeI
3856.02	SiII	4251.15	GaII
3862.59	SiII	4254.35	CrI
3933.66	CaII	4255.71	GaII
3950.35	YII	4261.92	CrII
3958.24	ZrII	4292.25	MnII
3982.59	YII	4314.08	ScII
3991.14	ZrII	4320.73	ScII
4005.71	VII	4325.01	ScII
4023.39	VII	4326.76	MnII
4030.76	MnI	4374.46	ScII
4033.07	MnI	4374.96	YII
4034.49	MnI	4383.55	FeI
4035.63	VII	4385.38	FeII
4067.05	NiII	4415.41	FeI
4077.71	SrII	4416.82	FeII
4111.01	CrII	4468.49	TiII
4122.64	FeII	4475.26	PII
4124.79	FeII	4481.2	MgII
4128.05	SiII	4539.62	CrII
4130.88	SiII	4539.79	CrI
4136.96	MnII	4540.50	
4149.22	ZrII	40.72	CrI
4163.64	TiII	41.07	
4171.90	TiII	4552.56	SiIII
4173.43	FeII	4554.03	BaII
4206.38	MnII	4555.08	CrII
4215.52	SrII	4555.89	FeII
4233.17	FeII	4558.66	CrII
4242.38	CrII	4563.76	TiII

TABLE III-3

ELECTRON PRESSURE

Star	Nm	B-V	θ	θ_{uv}	$(P_e)_{uv}$	$\Delta(\log P_e)_{uv}$
Normal						
ι Her	17	-.18	(.29)	(.37)	215	
τ Her	17+	-.15	(.31)	(.40)	175	
π Cet	18	-.14	.33	.42	120	
θ Aql	(20)	-.08	.39	.50	45	
μ Ser	18½	-.04	.43	.54	75	
ν Cap		-.01				
η Leo	25	(-.01)	(.46)	(.57)		
α Lyr	19½	.00	.47	.58	48	
α CMa	19^-	+.01	.48	.59	62	
O Peg	19½	+.01	.48	.59	47	
θ Leo	20^+	+.02	.49	.60	35	
β Ari	19^-	+.14	.58	.68	53	
72 Oph	19	+.16	.59	.69	49	
Peculiar						
HD34452	17	-.17	.30	.39	160	
108 Aqr	19:	-.14	.33	.42	100	
α^2 CVn	19^-	-.11	.36	.46	80	
HD172044	18^+	-.10	.37	.47	97	
4Cyg	18½:	-.10	.37	.47	87:	
112Her	18	-.09	.38	.48	104	
υ Her	19	-.09	.38	.48	70	-.10
HD3322	18½:	(-.08)	(.39)	.49	84:	
HD205087	18	-.08	.39	.49	87	
π^1 Boo	19^-	(-.08)	(.39)	(.49)	(76)	
HD224801	19	-.07	.40	.50	68	
HD192913	20^-	-.07	.40	.50	50	-.2
87Psc	19	-.07	.40	.50	67	
HD175744	19^-	-.05	.42	.53	70	
φ Her	19½	-.05	.42	.53	51	-.15
ι CrB	18^+	-.04	.43	.53	84	
17ComA	18	-.04	.43	.54	85	+.1*
HD10783	19^-	-.04	.43	.54	68	
ρ Her (A)	25	(-.03)	(.44)	(.55)	(33)	-.3
21Per	19^-	-.00	.47	.58	63	
HD8441	19	.00	.47	.58	58	
HD4778	19:	+.01	.46	.56	60:	
HD107162	19:	+.04	.51	.62	54:	
21Com	19	+.05	.52	.63	53	
HD2453	18^+	+.06	.53	.64	70	
52Her	19^-	+.08	.5	.64	58	

TABLE III-3 Cont.

Star	Nm	B-V	0	0_{uv}	$(P_e)_{uv}$	$\Delta(\log P_e)_{uv}$
HD204411	21	+.09	.54	.65	24	-.3
HD9996	19	+.10	.55	.66	53:	
9Tau	20^-	+.13	.57	.67	36	-.1
HD15144	19:	+.16	.59	.69	49:	
HD164258	19	+.16	.59	.69	49	
HD188041		+.20				
HD191742	20	+.22	.63	.71	33	
γ Equ	19^+:	(+.27)	.66	.73	42	

*Blanketing correction might bring it closer to normal

TABLES III-4 a to d

Symbols Used in Tables III-4 a to d

?   Information could not be obtained from the available material.

Information based on published sources.

*   Lines are not noticeable. At this temperature and for the dispersion used, normally, lines are not expected to be noticeable.

X   For the cases where the star has sharp lines and high dispersion was used, this means that lines are too weak to be detected.

$\not{S}$ Lines might be present but not clearly visible on the spectrograms available.

Blank Not investigated, or, lines are too wide to permit an estimate

:   Less reliable.

( )   Quite uncertain.

+   Strengthening (or overabundance).

-    Weakening (or underabundance).

S,VS Strong or very strong; lines not observed in normal stars.

Lines present but not strong; lines not observed in normal stars.

TABLE III-4a

Mn - Stars

	α And	HD172044	κ Cnc	μ Lep	υ Her	π' Boo	112Her	HD3322	87Psc	φ Her	ι CrB
B-V	-.10	-.11	-.11	-.10	-.09	(-.09)	-.09	?	-.07	-.05	$-.04_5$
ω (Å)	.8	.3:	.13	.3	.2:	.4	.15:	.5:	.5:	.3:	.07
Be	0	S	S	S	S	X	S			X	X
Mg	0	0:	-2	0	-2	0	?	+1:	?	-2	0
Si	+3	(0)	+3	+5	+6±	+3	+2	+3:	?	+3	0
P	*	?	S	Xor	X	X	S	?	?	X	X
Ca	0	0:	0	1	0	0	0:	0:	?	0	0
Sc	*	*	+5:	+6:	(+)	+7	*	0 or +	+6	+5	0 or 1
Ti	+3:	+4:	0:	+3:	+3	+3	+1:	+2:	+6	+2	0
V	*	0:	X	*	*	0:	(0)	(-)	(-4)	(-)	-5
Cr	(0)	-2:	(0)	(0)	-2	0	-2	-4:	-2:	+3	-1
Mn	+10	+9	+9	+10	+7	+9	+9	+10	+7	+6	+4
Fe	-3:	-2:	-2	-3	-3	-4	0:"	-6:	-2:	-6	-1
Co	*		X	*							
Ni	*	(-)	-2	-4	-4:	-6	(0)	-3:	-2:	-6	-4
Ga	*	?	S	S		S	S	?	?	X	X
Sr	0:	?	(+)	+5	+2	+2	?	(+1)	?	0	+5
Y	+9	+7	X	+8	+6	+7	+6	+7	+6:	+7	+6
Zr	*	+1:	(+)	*	*	0:	(0)	0:		+6	+3
Ba	*	?	X	*	(+2)	X	?	?	?	+1	0
Eu	*	?	X	*	X	X	?	?	?	X	X

TABLE III-4b

Si - Stars

Star	HD34452	108Aqr	τ^9Eri	HD124224	α^2CVn	4Cyg	HD205087	HD224801	HD192913	HD175744	21Per
B-V	-.17	-.15	-.12	-.11	-.11	-.10	-.08	-.08	-.07	$-.04_5$	-.00
Phase					EuII Max			EuII Strong			
ω (Å)	1:	.8	(.6):	(4)	.3	.4	.4	.8:	.2	1:	.6
Be		S	X		S		X		S		S
Mg	0:	-2	0	-	-3	-4	-6	-5	-3	-5:	-5
Si	+10	+7	+9	S	+7	+8	+7	+10	+6	+7	+7
P											
Ca	+3	-4	-6:		-3	-3	-5	-2	0	-7	-2
Sc	*	(+5)	*		+3	+3:	+6:	(+3)	+8	+8	-3
Ti	(+)	+2:	-3		+3	0:	-3:	+2	+3	(-5)	+3
V	*	*	*		(0)	*	+7:	(+1)	0:	*	-2
Cr	+10	+6	+7		+2 to 6[a]	+6	+10	+5	+5	+4	+6
Mn	(0)	0:	(+3)	(0)	+4	0:	+2	+3	+4	(0 or -)	+6
Fe	+5:	+5:	+4:		+3	+4:	(+5)	+4:	+5	+3:	+3
Co	*									*	*
Ni	*	-2	-3:		-2	(-)	-4:	-3:	-5	*	-6:
Ga									X		
Sr	(+)	+5:	+5		+5	+2	+7	+6	+7	?	+7
Y	*	*	*		+3:	*	+3	+2:	+3:	*	+7
Zr	*	*	*		+3	(+)	+7	0:	+6:	*	+2:
Ba	*	*	*		-3	*	*	*	0 or +	?	0:
Eu	∅	*?	∅		VS	∅	S	S	VS	?	S

Remarks: a. : - value at CrII max. phase.

TABLE III-4c

Cr - Stars

Star	17ComA	HD10783	HD125248	HD4778	HD153882	γAri(S)	HD2453	HD9996	9Tau	HD188041	γEqu
B-V	$-.04_5$	$-.04_5$	+.01	+.02	+.04	?	$+.07_5$	+.10	+.13	+.20	+.27
Phase			EuII Max.							EuII Max.	
ω(Å)	.4:	.3:	.18	1/2-1	.4:	.8to 1	.14	.1	.15	.11	.09
Be											
Mg	0:	(-1)	-3	-4	0	0:	0:	0	0:	0	+2
Si	(-)	(+1)	-6	(0)	-4	0:	+3:	-2:	+4:	(-)	+3
P											
Ca	-7	-9:	-6	-7	-9	-6:	-6:	-8	-7	(-4)	(-4)
Sc	(0)	-3:	+7	(0)	+1		0	+4	+1	0:	-3
Ti	-2	+1	-1	+3	-6		+6	-3	0:	0:	0
V	-:	-3	(0)	(-1)	-6		0	-6	-3:	-3	0
Cr	+9	+9	+5 to +9	+6	+11	+9	+11	+7	+6	+9	+4
Mn	+6	+3	+6	+6	+2	(+4)	+6	0:	-1	+4:	0:
Fe	0:	-3:	+6	0:	+6	+3:	+4	0	0	+2	+4
Co							(0)		(0)		(+6)
Ni	(-)	-4	-4	-4:	-4:		(0)	0:	-3	-6	0 or +
Ga											
Sr	+5	+4	+3	0	0	+5:	+5	+4	+2	+6	+7
Y	(+1)	+3	+2	-1	-7		0:	+1:	-1	+1	+2
Zr	(0)	(-)	+6	(0)	-6	-6:	0:	-4	-1:	+6	+1
Ba	?	?	-9	-6	-10		(-6)	-9	?	-6	-1
Eu	Poss.	(+6)	+9	+8:	X	+9	+9	X	?	+12	+9

Remarks: b. Value at CrII max. phase.

TABLE III-4d

Sr - Stars

	HD8441	HD107612	21Com	52Her	HD15144	HD164258	HD191742
B-V	0.00	+.04	+.05	+.08	+.15	+.16	+.22
ω (Å)	.08	.8:	1:	.4	.15	1:	.12
Be							
Mg	-3	?	2	0	+3	0:	?
Si	-7	-6	(-)	(-4)	0	(-)	(-)
P							
Ca	-10	-1	+3	-3:	+3	-6	-1
Sc	+4	+4	(+2)	+2	+6	(+)	-2
Ti	-4	+1	+5:	-1	+4	-1:	+4
V	-6	-3:	(0)	-3	-6	-6:	+1:
Cr	+10	+6	+5	+9	+7	+9	+6
Mn	0	(-3)	(+4)	+4	(+4)	0	0
Fe	-1	(+1)	(-1)	+4	+3	0:	0
Co	(0)					+:	+:
Ni	-6	(0)	(0)		(+)	-4:	0:
Ga							
Sr	+10	+12	+11	+7	+9	+8	+10
Y	-4	(-1)	(-1)	0	0	0:	(0)
Zr	+3	(+)	(+1)	+1	+6	+6	+5
Ba	-9	(0)	?	-3	-2:	?	?
Eu	+5	?	?	(+8)	+8	Possible	∅

TABLE III-5

	α^2 CVn		κ Cnc	
	Curve of Growth (*)	Estimates	Curve of Growth (#)	Estimates
Mg	-.4	-3	-.8	-2
Si	+1.0	+7		
Ca	-1.7	-3		
Sc	-.2	+3		
Ti	+.4	+3	-.2	0:
V	+.1	(0)		
Cr	+.7	2 to 6	+.1	(0)
Mn	+1.2	+4	+1.5	+9
Fe	+.5	+3	-.3	-2
Co				
Ni	+.5	-2		
Sr	+1.1	+5	+1.5	(+)
Y	+1.3	+3:		
Zr	+1.5	+3		
Ba	$\lesssim$0.0	-3		

* Burbidge and Burbidge (20)

Sargent (65)

TABLE III-6

i	$\log\left(\frac{\epsilon_i}{\epsilon_H}\right)_{\alpha\, CMa} - \log\left(\frac{\epsilon_i}{\epsilon_H}\right)_{\alpha\, Lyr}$	
	Boyarchuk and Hunger	Estimated
H	0	
He	+.3	
Be	<-1.5	
Na	+.5	
Mg	+.6	0 for MgI, +.3 for MgII
Al	+1.4	+.5
Si	+1.6	+.5 or +.3
Ca	-.5	-.3
Sc	-.2	+.2
Ti	+.6	+.5
V	+1.0	?
Cr	+1.0	+1.0
Mn	+.7	+.4
Fe	+.9	+.8
Ni	+.5	+.6
Sr	+.4	+.6
Y	+1.1	+.5
Zr	+.2	<+.3
Ba		+1.0:

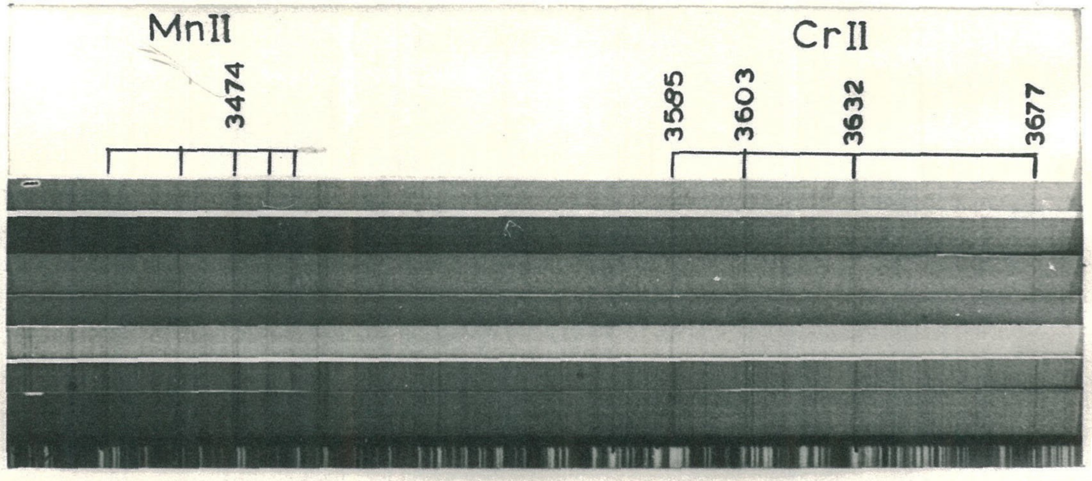
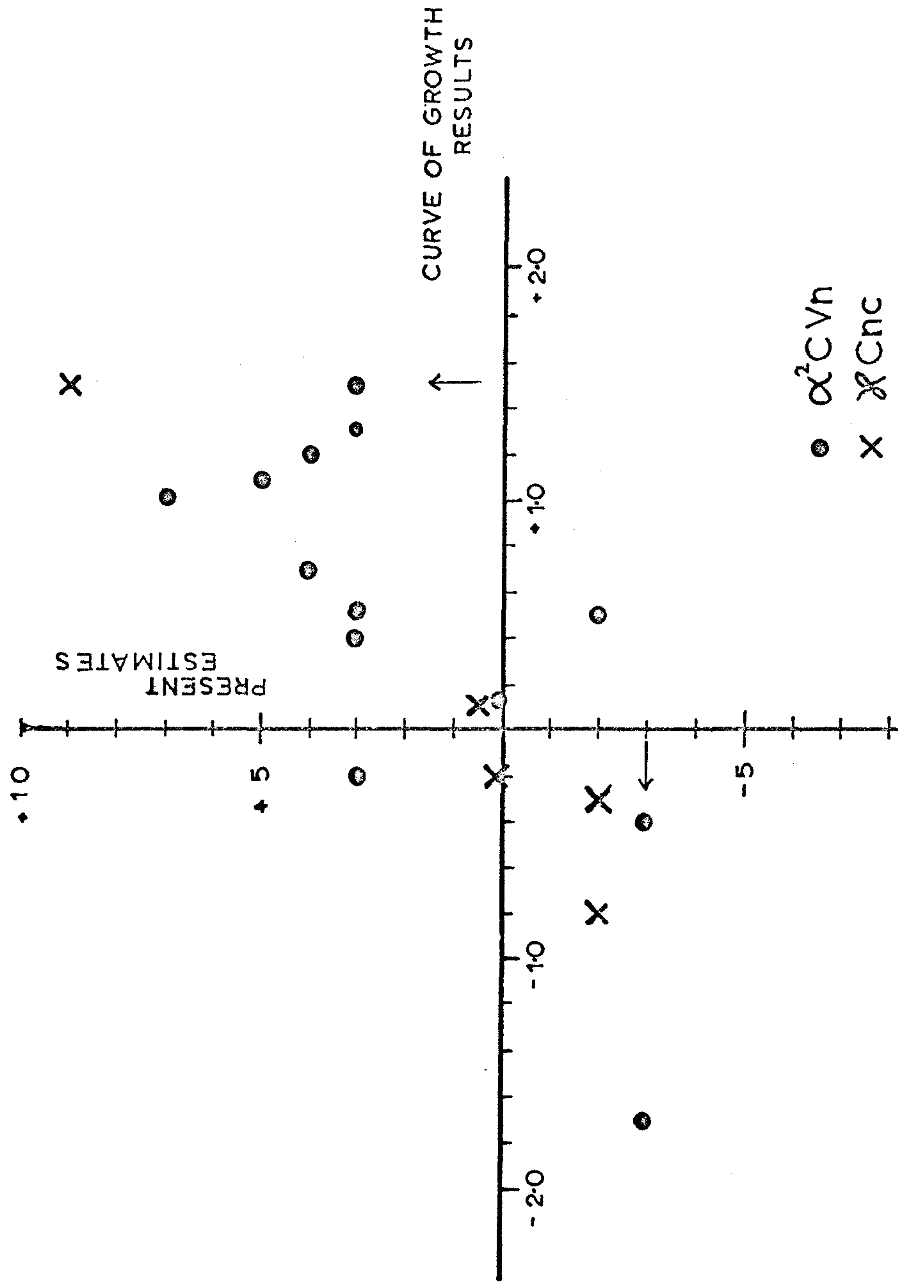


Figure III-1
Showing easily recognizable patterns of CrII and MnII in the ultraviolet.


CURVE OF GROWTH RESULTS
PRESENT ESTIMATES
+2.0
+1.0
-1.0
-2.0
+10
+5
-5
-10
α²CVn
χ Cnc

'igure III-2

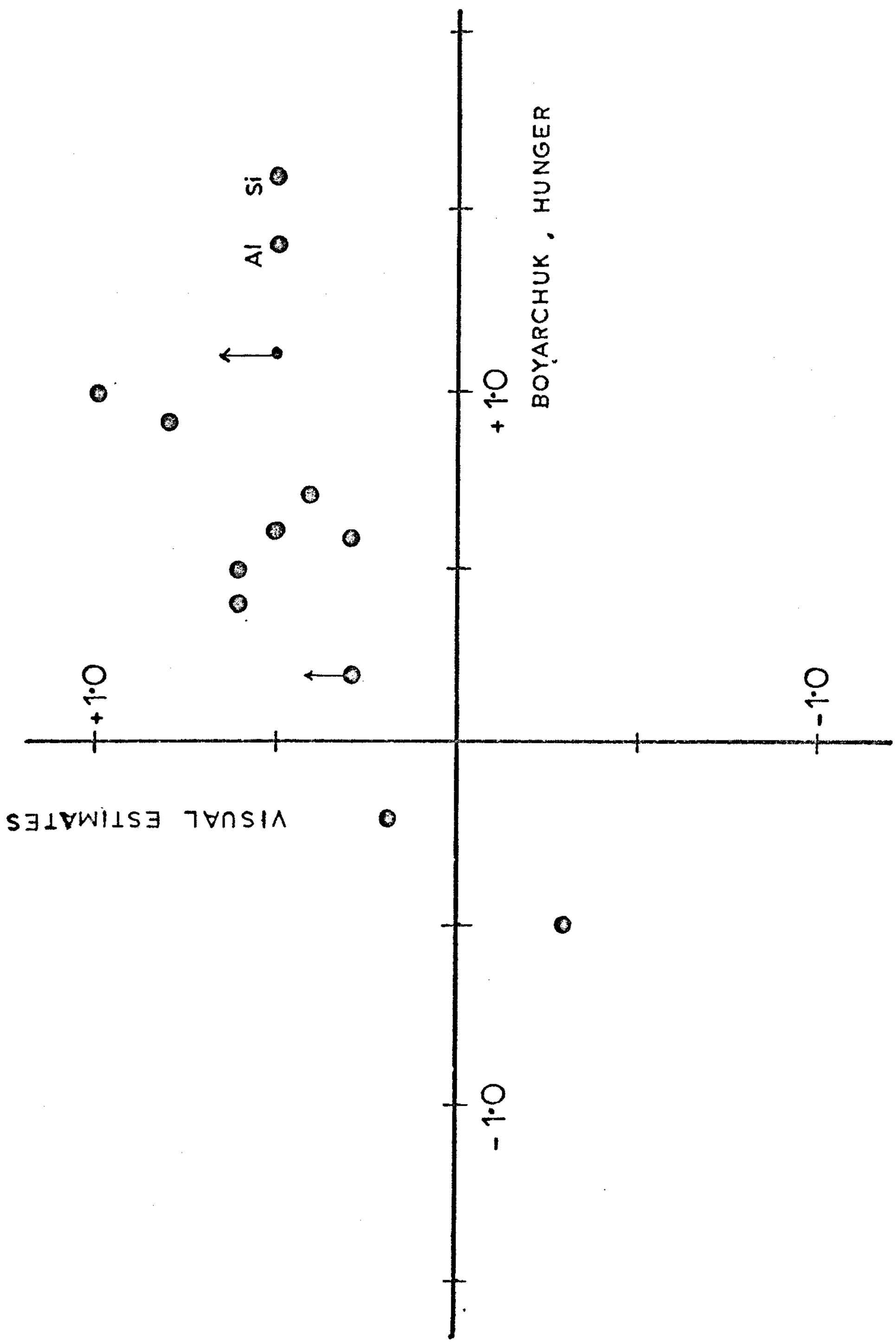


Figure III-3

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CHAPTER IV

THE RARE LIGHT-ELEMENTS

Sec. IV-A

Observability of D, Li, Be, and B

For study of the mechanism of nucleosynthesis in stars, abundance determination of the rare light-elements - deuterium, lithium, beryllium, and boron - is of special significance. Total terrestrial and meteoritic abundance of Li, Be, and B is only about one-millionth that of C, N, and O; under the physical conditions prevailing in the stellar interiors, these light elements are rapidly destroyed by thermonuclear reactions with protons. But, in certain Ap stars, some of these light-elements, as well as some other heavier elements, have greatly strengthened spectral lines presumably due to an overabundance of these elements by large factors.

First we review the present status of the observability of these rare light elements:

Deuterium - Because of the high degree of stark broadening and small separation between the deuterium and hydrogen lines, the $\frac{D}{H}$ ratio cannot be observed to an accuracy much better than .01 (66). The terrestrial value for this ratio is about 10^{-4} (35). No positive observations of deuterium have been made in astronomical objects;

and observationally, the upper limit set for this ratio, by present methods, is 10^{-2} to 10^{-3} . Consequently, the chances of observing deuterium in Ap stars are small.

Lithium - The ionization potential of LiI is 5.4eV and that of LiII is 75.6eV. Therefore, in the atmospheres of the Ap stars, all the lithium exists in the LiII state. Ground-level lines of LiII lie in the unobservable part of the ultraviolet. Excitation potential of the observable lines is about 60eV. Therefore, detection of LiII in Ap stars, with the ground-based telescopes, in not feasible.

Beryllium - Of the rare-light elements, beryllium is most favorable for spectroscopic detection in the Ap stars. In fact, this was a main motivation for undertaking the present study of the ultraviolet spectrum of the Ap stars down to 3100Å . Results on beryllium are discussed in Section B and the ensuing sections of this chapter.

Boron - Ionization potentials of BI and BII are 8.3eV and 25.1eV respectively, and, therefore, boron is expected to be found, in Ap stars, only in the neutral or singly ionized states. However, no low-excitation lines of BI, BII, or BIII lie in the observable region of the spectrum. Only two BII lines, with high excitation potentials, are listed in the RMT: λ3451.4 ($2\,^1P^\circ - 2p^2\,^1D$; 9.06eV) and λ4121.95 ($3^3D - 4^3F^\circ$; 18.60eV).

Section IV-B

Survey of Beryllium

Earlier Work. Greenstein and Tandberg-Hanssen (38) have studied BeI triplet ($\lambda\lambda$3321.01, -.09, and -.35) and the BeII double ($\lambda\lambda$3130.42, -31.06) in the solar spectrum, and have derived for the abundances of H and Be, by number, the logarithmic ration, $\left[\frac{H}{Be}\right] = 10$. Of these lines, λ3131.064 and 3321.013 - .086 are quite blend-free and reliable, and have equivalent widths 79.4mA and 6.6mA respectively.

Bonsack (17) has studied abundance of beryllium in four bright stars - α CMa (A1V) , γ GemA (A1V), α Lyr (AOV), and the Ap star α^2 CVn (AOp) - by estimating the contribution by the BeII 3130.42 to the underlying VII line λ3130.26. The equivalent width of the contribution due to BeII 3130.42, and the total equivalent width of the blend (in parentheses), in mA , are $<$4(56), 8(26), and 64(91) and the derived logarithmic ratio $\left[\frac{H}{Be}\right]$, $>$ 12.1 , 10.6 , and 8.8 for α CMa, α Lyr, and α^2 CVn respectively.

Sargent, Searle, and Jugaku (68) have made a survey of BeII 3130.31 doublet in 25 Ap stars. In this abstract, they note "Of the ten Mn stars surveyed, four (χ Cancri, 112 Herculis, μ Leporis, and υ Herculis) have strong BeII lines; hence we conclude a Be abundance about 100 times that in the sun. The BeII lines are absent, or weak or very weak in five stars showing the λ4200 line of SiII and in the remaining six Mn stars." Names of the Si-4200 stars, of their survey, are not given.

<u>Present Work</u>. The first and second ionization potentials of beryllium are 9.3eV and 18.2eV . For log Pe = 2 and 3 , the range of the Ap stars, the number of neutral atoms equals the number of singly ionized ones at θ = .81 and .72 respectively. Therefore, in the atmosphere of the Ap stars, both Be and Be^+ states are expected to be present. Strongest lines of BeII are the zero-volt resonance doublet lines at 3130-31 . But, for BeI the ground-level line, λ2348.6 is unobservable; the strongest lines in the observable region are those of the 3321 triplet ($2^3P^\circ - 3^3S$) arising from the 2.71eV level. These triplet lines of BeI have very low f-values (.034 as against .50 and .25 for the doublet lines of BeII) and are very much weaker than the BeII doublet even in the solar spectrum. Population of Be atoms in the lower levels of the BeI triplet (2.71eV) relative to the population of Be atoms in the ground level of BeII increases, in the beginning, as we go to temperatures hotter than the sun. However, before the temperature of the Ap stars is reached, the relative population of BeI

2.7eV level is again very low. Consequently, only the BeII doublet λ3130-31 is suitable for detection of Be in the Ap stars. (Other lines of BeII, listed in the RMT, originate from 12eV levels and therefore are unfavorable.)

Blends with the BeII 3130-31 Doublet

At the dispersion of 21Å/mm, used for this survey, lines of this BeII doublet and the underlying normal lines are not resolved. Therefore, it is important to investigate what are the possible lines that might make the 3130-31 blend strong. Of the lines listed in the MIT tables, only those in Table IV-1 need to be investigated as candidates which might contribute to the λ3130-31 blend in the Ap stars.

Out of the lines listed in this table, only VII 3130.27, FeII 3130.57, TiII 3130.80, and the BeII doublet make any significant contribution in the normal A-type stars such as αLyr. The Cr lines might contribute in stars with very large overabundance of chromium. The EuII lines will make a contribution in the cooler Ap stars with very strong EuII lines (e.g. γEqu. or HD188041). Likewise, a contribution by the YII line will be significant in the cooler Ap stars only if yttrium is overabundant by a large factor. Lines of PII have been observed only in the Ap stars κCnc and 112 Her. Inspection of PII 4475.26 (I = 150) suggests that in κCnc, PII 3130.38

will make a minor, probably non-negligible, contribution. The contribution by the silicon line $\lambda 3130.40$ will be unimportant: In τ^9 Eri which has very strong silicon lines, the equivalent width of the total feature at $\lambda 3130.5$ is less than 25mÅ. In other types of Ap stars, therefore, contribution by this silicon line, if any, will be negligible.

The columbium (or niobium) line at $\lambda 3130.79$ is among the stronger lines of CbII as listed in the RMT. The strongest of the CbII lines lie in the $\lambda 3100$ region, and have been noticed in the solar spectrum. But this region has not been adequately observed in other types of stars. To ascertain whether the contribution, if any, due to CbII 3130.78 is significant, $\lambda\lambda 3094.17$, 3145.40, 3163.40, 3194.98, 3206.35, and 3225.48 of CbII were examined in those stars in which $\lambda 3130.31$ blend is abnormally strong for their B-V value. On the basis of this examination, it was found that lines of CbII might be present in υ Her, HD192913, and HD9996, but the contribution to 3130-31 is insignificant except in HD9996 in which the contribution may be minor.

Result of the Beryllium Survey: Following are the results of a survey of 25 Ap stars on whose spectrograms $\lambda 3130$ region can be seen.

(a) Stars in which $\lambda 3130.31$ blend is strong: 108Aqr, κCnc, α^2CVn, HD172044, υ Her, HD192913, 21 Per, and HD9996(?). Further discussion of these stars and an analysis of the beryllium abundance is given in the next section. (Earlier, Sargent et al found strong BeII lines only in Mn-stars: κCnc, μLep, 112 Her, and υHer). Attention is drawn to the Si-4200 stars in which we find strong Be.

(b) Stars in which BeII lines make no significant contribution to the λ3130.31 blend: τ^9Eri, HD205087, 17 ComA, 87 Psc, HD175744, ϕHer, HD10783, π^1Boo, HD2453, 52Her, and 9Tau.

(c) Stars in which BeII contribution to λ3130.31, if any, is small: 4Cyg, HD107612, 21 Com, HD15144, HD164258, and γEqu.

The above results (a) and (b) were derived from very well exposed plates. But, for stars in (c), spectrograms were somewhat less strongly exposed in the λ3130 region; consequently, a possibility of a weak contribution by the BeII doublet cannot be ruled out. Figure IV-1 shows tracings of 10 A/mm plates of αLyr and HD192913 in the region of the BeII doublet. In HD192913, BeII lines are very strong. Figure IV-2a shows tracings of 21A/mm plates in which at the position of the BeII doublet a strong blend is present. Of these, all, except HD9996, have BeII as the major contribution to the blend. A similar Figure, IV-2b, shows Mn, Si-, and normal stars in which BeII lines are not strong.

Section IV-C

Beryllium Abundance Determination

Equivalent Widths: The equivalent width of the λ3130-31 blend in the eight Ap stars of group (a) and six normal stars, based on 21 A/mm plates, is given in Column 3 of Table IV-2; for HD192913 and α Lyr, measurements are based on 10 A/mm plates. Column 4 gives an interpolated (or extrapolated) equivalent width of this blend for a normal star with the same B-V value as the Ap star. The next columns give the strengthening or weakening of V, Ti, and Fe - main contributors to this blend in the normal stars - in each of these Ap stars; this information has been taken from the Tables III-4 of the previous chapter. Column 9 lists the other significant contributors, besides BeII, in each of these stars. On the basis of all this information, then, in Column 10 we give the estimated contribution by the BeII lines λλ3130.42, -31.06. These BeII lines are formed by transitions $2\ ^2S_{1/2} - 2\ ^2P^{\circ}_{3/2}$ (3130.42) and $2\ ^2S_{1/2} - 2\ ^2P^{\circ}_{1/2}$ (3131.06), and should have f-values in the ratio 2:1. If these lines are on the flat part of the curve of growth, we should expect that λ3130.42 contributes close to 50 percent of the BeII contribution; on the other hand, if the BeII lines are weak, we should expect a 67 percent contribution. The true value should be somewhere in between. However, abundance analysis is made for both these sets of values; this will provide some estimate of error.

Curve of Growth Analysis: The accuracy of the equivalent widths derived in the above manner - which is the best that can be done unless higher dispersion plates are available - does not call for a very precise analysis, therefore, our aim is to make an order of magnitude determination of Be abundance.

The abundance analysis is made using a theoretical curve of growth based on the Milne-Eddington model (Wrubel 78). Bonsack (17) has determined the Be abundance in α^2CVn relative to the sun. In the present work we shall investigate the beryllium abundance in these stars relative to α^2CVn. Using the usual notation, (Aller (3))

$$\log \frac{N_r}{N_{r_c}} = \log \frac{\eta_o}{\eta_{o_c}} + \chi_{r,s} (\theta - \theta_c) + \log \frac{\varkappa_\lambda}{\varkappa_{\lambda_c}} \cdot \frac{v}{v_c} \frac{u_r(\theta)}{u_r(\theta_c)}$$

gives the relation for each line observed in a given star and a comparison star; quantities for the latter are denoted by a subscript "C". This relation is here applied to BeII line λ3130.42 (χ = OeV). The various quantities required for this formula are derived in the following manner.

(i) Electron Pressure: The value of electron pressure is needed in the Saha-ionization formula and for the determination of the coefficient of continuous absorption, κ_λ , in the region where the absorption line is formed. The value of electron pressure, for each star, is derived by using the Inglis-Teller formula, and θ_{uv} (defined later) and n_m , the quantum number of the lower level of the Balmer line at the confluence point. Usually, this method gives a value of Pe which is much smaller than that derived from the wings (formed in the deeper layers of the atmosphere) of the earlier members of the Balmer series. For AOV stars, according to the usually quoted values, log Pe = +3, while from the Inglis-Teller formula we get values around +2. Therefore, the use of the present method needs justification:

In the A-type stars, the coefficient of continuous absorption in the ultraviolet is about an order of magnitude greater than in the blue region. The effect of this increased absorption in the UV is to raise the effective level of line formation to a higher and cooler region. The electron pressure in this region will, naturally, have a smaller value. Therefore, this smaller value of electron pressure, derived from the confluence point of the Balmer series, is more appropriate for the lines in the UV than the higher value of Pe existing in the deeper layers of the atmosphere.

(ii) Temperature: Temperatures for these stars have been derived from their B-V color in a manner discussed in Section II-C. (In that section, we have already discussed applicability of the (B-V)-temperature relationship.) Temperatures derived from B-V color are usually closer to the ionization temperature than to the excitation temperature. Since $\chi = 0eV$ for the line under consideration, and since the partition function for BeII varies very slowly for the range of temperatures encountered here [(Aller, Table 8A), (3)], we need only the ionization temperature for calculating the fraction of Be atoms in the singly ionized state.

The line under investigation lies on the short wavelength side of the Balmer discontinuity and, as discussed earlier, is formed at a relatively higher and cooler layer due to the increase in continuous absorption. For this reason, the values of θ_e dervied from B-V colors have been converted into corresponding 0_{uv} values in the following manner: It was assumed that, effectively, lines are formed at an optical depth $\tau_\lambda = .6$. Then, from Mihalas' models of stellar atmospheres for log g = 4 (main sequence) and different values of θ_e, the value of $\theta_{(\tau_{3646-} = .6)}$ - which we call here θ_{uv} - was noted. In Table IV-3, both the θ_e and θ_{uv} values have been listed.

(iii) The velocity of small scale motions, v, (which includes microturbulence and thermal motions) is taken to be 4 km/sec for all the stars. Ideally, v should be determined individually for each star. However, since (a) it is not certain that in these peculiar objects turbulent velocity derived from lines of some other elements will be applicable to the BeII lines, and (b) the equivalent width contribution by BeII lines is only approximately known, a uniform value of v will suffice for our present purpose. Earlier investigators (see Deutsch (29) and Bonsack (17)) have derived, for normal and peculiar A-type stars, values of v in the range 1.5 to 4.5 km/sec; values around 4 km/sec being more typical for the Ap stars.

(iv) Continuous absorption, $\varkappa_\lambda$ (λ = 3200Å), values have been taken from the Mihalas tables (53).

Calculations and results are presented in Tables IV-3 and IV-4. In α^2CVn, Bonsack has derived for Be a logarithmic overabundance of +1.8 relative to the sun. In Table IV-4 values relative to the sun are obtained using this value.

Accuracy of Results: Differences between (a) and (b) values in Table IV-4 reflect the effect of increasing equivalent widths in (a) by about 33 percent. On this basis it is estimated that the results are probably within ±.5 of the true value, though, in some cases errors may be larger. Results for 108 Aqr and HD9996 may, in particular, have larger errors - the former due to uncertainty in temperature and the latter, due to the uncertainty of blends (uncertainty in θ becomes serious for $\theta < .4$ where a small error in θ sharply alters the fraction of Be atoms in the singly ionized state). Nevertheless, factors around 500 to 1000 for HD192913 and HD172044 are reasonably certain. All these stars (except 108 Aqr with $w = .8\text{Å}$) have quite sharp lines and, therefore, by observing them with 5 Å/mm dispersion, or, with sufficiently widened spectrum at 10 Å/mm, the accuracy of the results can be improved considerably.

TABLE IV-1

Lines Near the Beryllium Doublet

λ (Å)		Intensity RMT	Intensity MIT Arc	Intensity MIT Spark
3131.54	CrII	5		
31.07	BeII	30	200	150
30.80	TiII	15	20	100
30.79	CbII	1500	100	100
30.74	EuII	80	100W	100
30.57	FeII	2	4	4
30.57	Cr		1	12
30.48	Si		-	5
30.42	BeII	50	200	200
30.38	PII		-	[30]
30.27	VII	100	50	200
29.93	YII	40	8	50

TABLE IV-2

	B-V	3130-31 W(Å)		Blends					Be Contribution W(Å)	W_{BeII} 3130.4 (mÅ)	
		In Ap Star	In Normal Star	Ti	V	Fe	Cr	Other		(a) 50%	(b) 67%
(1)	(2)	(3)	(4)	(5)	(6)	(7)	(8)	(9)	(10)	(11)	(12)
Ap Stars											
108 Aqr	-.15	.35	.02	+2:		+5:	+6	Si	.25	125	165
HD172044	-.12	.28	.04	+4:	0:	-2:	-2:		.20	100	135
κ Cnc	-.11	.30:	.04	0:		-2	0	PII	.15	75	100
α² CVn	-.11	.158	.04	+3	(0)	+3	+4	Si	.10	50	67
υ Her	-.09	.27	.05	+3		-3	-2		.20	100	135
HD192913	-.07	.40*	.07	+3	0:	+5	+5	Si	.25	125	165
21 Per	.00	.26	.10	+3	-2	+3	+6	Si	.15	75	100
HD9996	+.10	.29	.16	-3	-6	0	+7	CbII(?)	.15	75	100
Normal Stars											
ν Cap	-.05	.15									
μ Ser	-.04	.09									
α Lyr	.00	.094*									
α CMa	+.01	.12									
72 Oph	+.16	.20									
21 LMi	+.17	.17									

*Measured on spectrograms with 10 A/mm dispersion

TABLE IV-3

Star	θ_e	θ_{uv}	$\log P_e$	$\log \kappa_{uv}$	$\log \frac{W}{b}$ (a)	$\log \frac{W}{b}$ (b)	$\log \eta_o$ (a)	$\log \eta_o$ (b)
108 Aqr	.33	.42	2.0	.96	.47	.60	2.07	2.82
HD172044	.35	.45	2.0	1.12	.38	.50	1.67	2.20
κ Cnc	.36	.46	1.9	1.07	.25	.37	1.23	1.63
α^2CVn	.36	.46	1.9	1.07	.08	.21	.81	1.14
υ Her	.38	.48	1.9	1.17	.38	.50	1.67	2.20
HD192913	.40	.50	1.75	1.17	.47	.60	2.07	2.82
21 Per	.47	.58	1.85	1.15	.25	.37	1.23	1.63
HD9996	.55	.66	1.7	.45	.25	.37	1.23	1.63

TABLE IV-4

Star	$\left[\frac{Be}{H}\right]_{Star} - \left[\frac{Be}{H}\right]_C$			
	$C \equiv \alpha^2 CVn$		$C \equiv$ Sun	
	(a)	(b)	(a)	(b)
108 Aqr	1.4:	1.9:	3.2:	3.7:
HD172044	.9	1.2	2.7	3.0
χ Cnc	.4	.5	2.2	2.3
α^2 CVn	.0	.0	1.8	1.8
υ Her	.9	1.1	2.7	2.9
HD192913	1.3	1.7	3.1	3.5
21 Per	.4	.5	2.2	2.3
HD9996	(-.3)	(-.2)	(1.5)	(1.6)

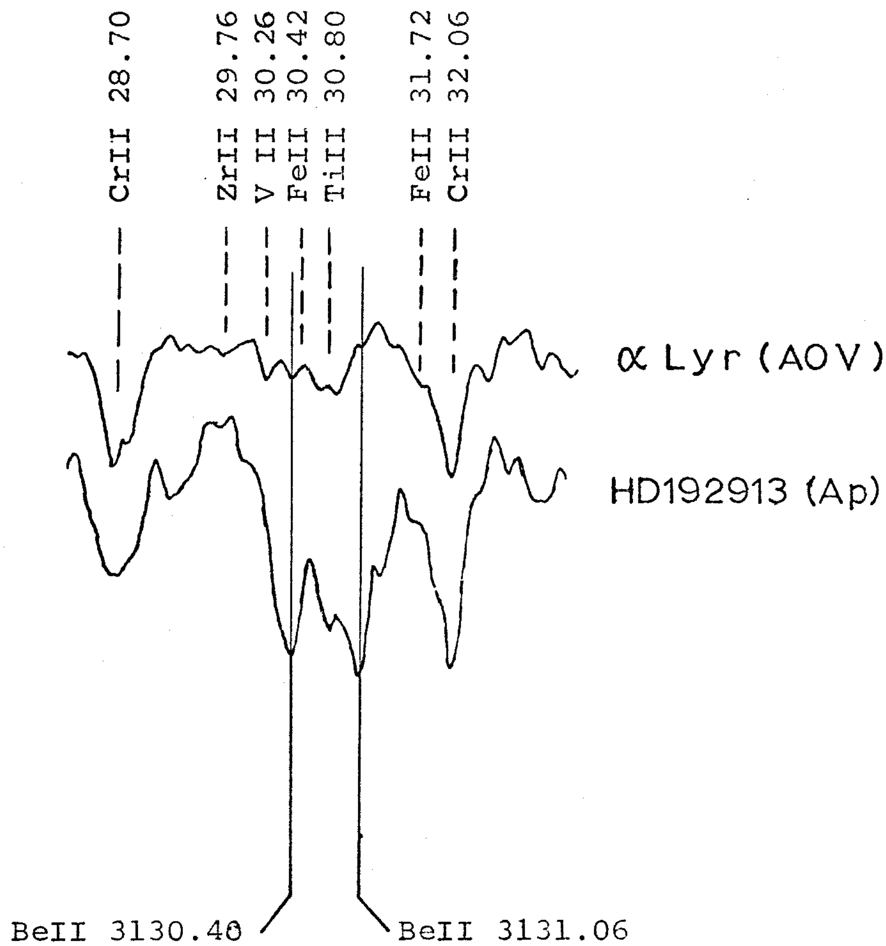
CrII 28.70
ZrII 29.76
V II 30.26
FeII 30.42
TiII 30.80
FeII 31.72
CrII 32.06
α Lyr (AOV)
HD192913 (Ap)
BeII 3130.40
BeII 3131.06

Figure IV-1

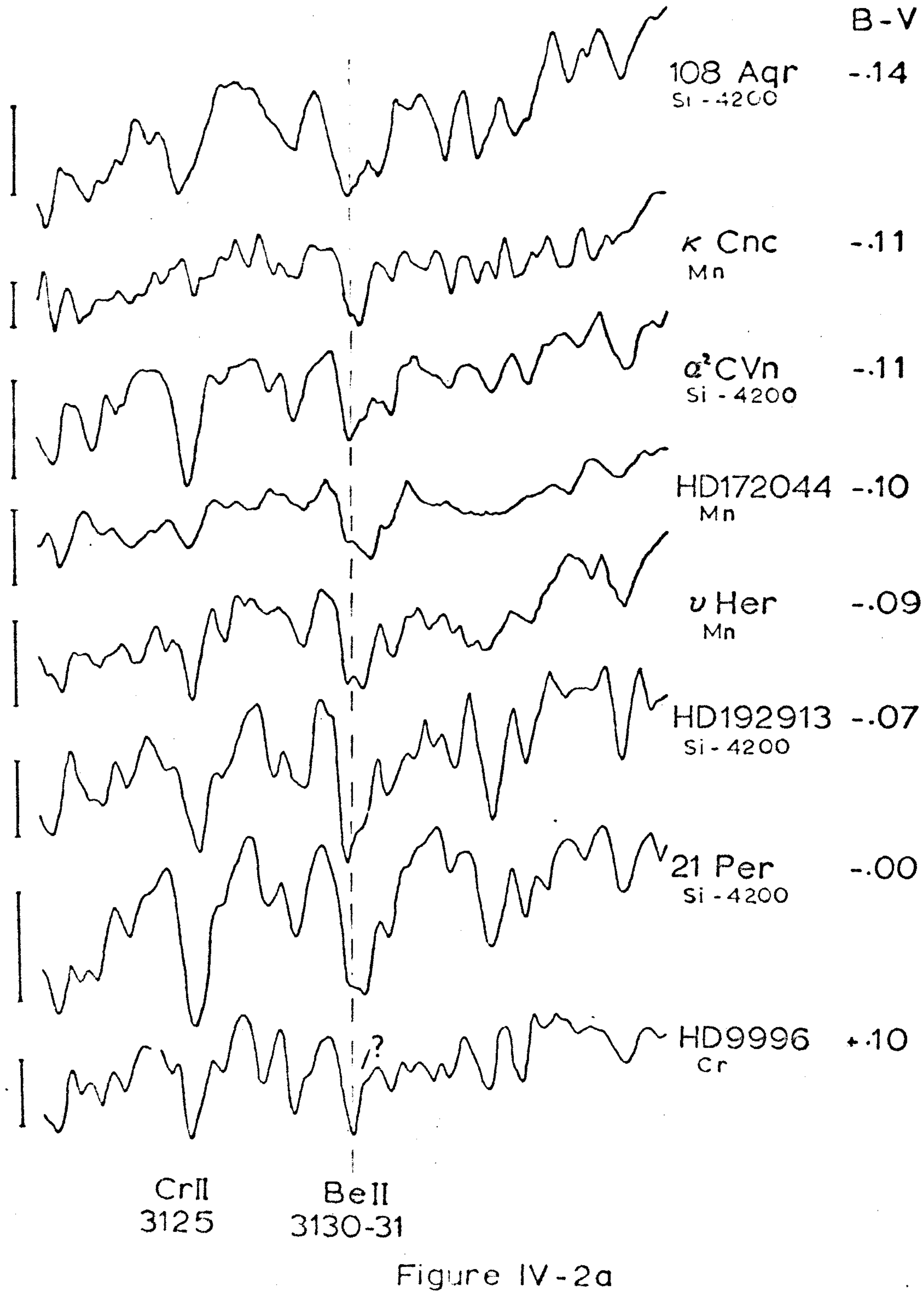

Figure IV-2a

In Figs IV-2 a & b the vertical lines represent 20% of the continuum

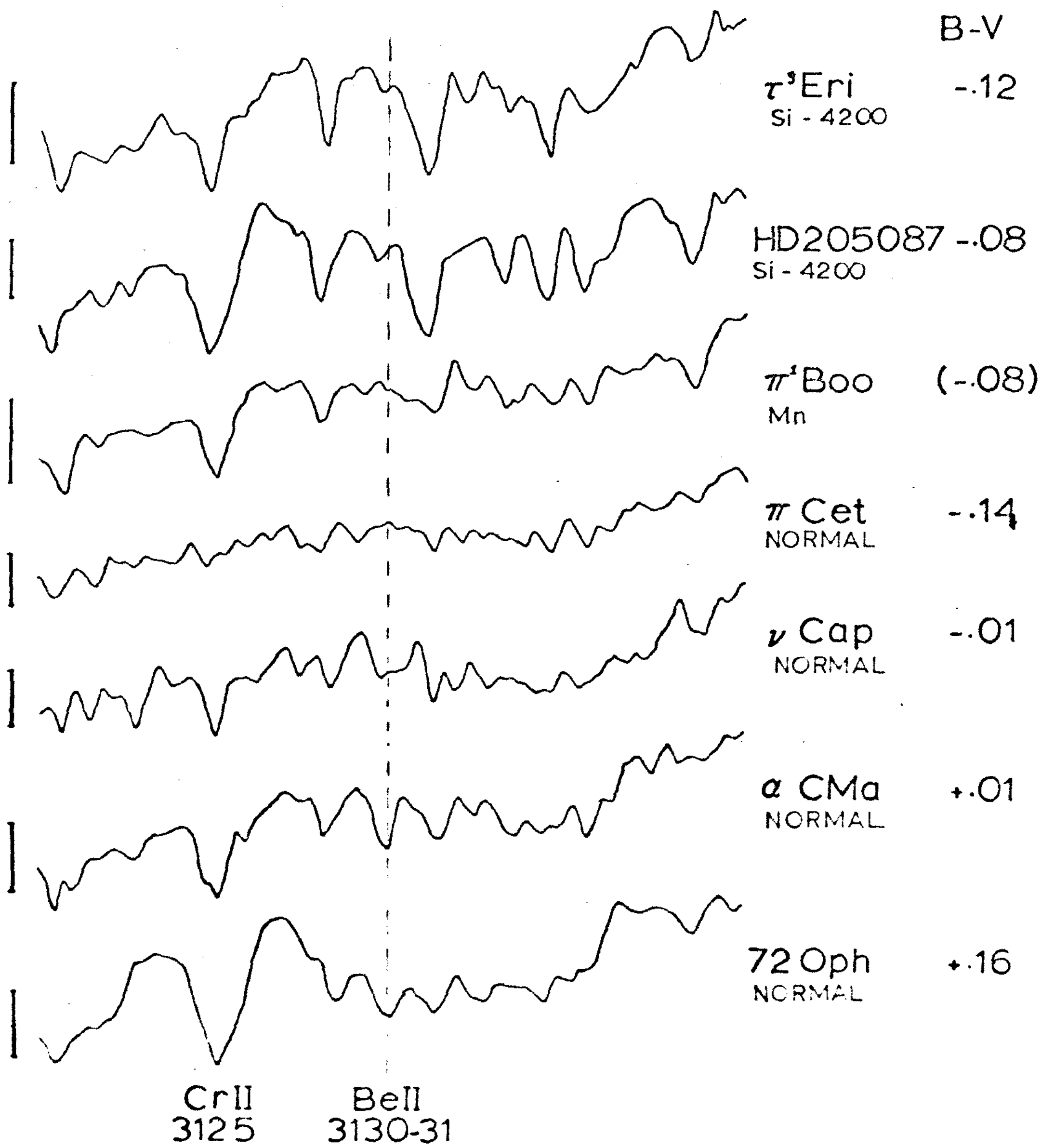
B-V
τ^3Eri
Si - 4200
-.12
HD205087 -.08
Si - 4200
π^1Boo
Mn
(-.08)
π Cet
NORMAL
-.14
ν Cap
NORMAL
-.01
α CMa
NORMAL
+.01
72 Oph
NORMAL
+.16
CrII
3125
BeII
3130-31

Figure IV-2b

CHAPTER V

The Rare-Earth Elements

The rare-earth elements, which form a group of fifteen elements from Lanthanum to Lutecium (Z = 57 to 71), are particularly associated with the Ap stars. Lines of EuII, which undergo a spectacular variation in intensity within a period of 5.6 days in α^2CVn were first identified by Baxandall (10). Of all the rare-earths, lines of europium are found to be the strongest and have been observed in many Ap stars (however in some cases, such as HD204411, stars designated "Eu-Cr" on the basis of low dispersion work might not have, in actuality, any strengthening of Eu lines). Lines of other rare-earth elements are relatively much weaker, and information about them has been recorded infrequently; in his Catalogue of Magnetic Stars, Babcock has made sporadic comments on the presence of Dy, Gd, or Ho lines in some of the stars. Most extensive work on the identification of the rare-earth elements has been done by Struve and Swings (73) and by Swings (74) for α^2CVn. But, most of the lines in the spectrum of this star are appreciably blended. Consequently, Struve and Swings have listed several contributors for each of the observed lines. These identifications have remained the basis for subsequent work on the rare-earths in other stars. (In subsequent work, however, not all the rare-earths have been investigated; and, some of these investigations are based on spectrograms with insufficient dispersion.)

The purpose of the present investigation then, is,

(a) To establish a group of "test lines" which are unblended, safe, and suitable for detection and for measurement of each of the rare-earth elements.

(b) To discover the rare-earth elements in other Ap stars, and

(c) To study differences in the pattern of distribution of various rare-earth elements in different Ap stars.

Section V-A

Identification Criteria

Stars in which rare-earth elements are present are usually found to be crowded with spectral lines. Therefore, for a reliable survey of all the rare-earths, the use of high dispersion becomes necessary; only the high-dispersion plates (10 Å/mm) were used for this work. "Test-lines" have been selected for use with high dispersion (and for stars with not too wide lines). In other words, those rare-earth lines have been selected which are free of blends to within .5 Å on either side.

By the use of the RMT, the MIT Tables, and the Tables of Spectral-Line Intensities (50), about 15-20 strongest lines for each element (singly ionized) were selected. These lines were tested against blends by locating the position of these lines on a tracing of the AIV star θ Peg. Next, with the help of the above mentioned tables, these lines were checked against possible blends by lines of those elements - including the rare earths, FeI, FeII, etc., - which are abnormally strong in the different types of Ap stars. Special care was taken with the chromium lines because these are found to constitute a major source of unexpected blends. Lines which passed this screening procedure are listed in Table V-1. Nearly all these lines can be used reliably in most cases. However, if some element is unusually overabundant, it might still make a significant contribution as a blend. Therefore, to take their effects into account, all such contributors (which might become significant in an Ap star of one type or the other) have been listed against each line in Table V-1. It is to be emphasized that, (a) none of these "Possible Contributors" will be significant in most cases, and (b) the "Possible Contributors" do not include all the possible lines in the immediate vicinity of a given line; under this column only those lines are listed which have some likelihood of being present, and which should be verified in the cases where presence of blending lines is suspected because of an inconsistency of the line strength with the other lines of the same element.

Laboratory data for most of the rare-earth elements is scant, therefore selection of "15-20 strongest lines" becomes difficult. Even the ionization potentials of several of these atoms and ions are not available. The intensity values listed in the above-mentioned three tables are often discordant because these laboratory intensities represent different excitation conditions. In such circumstances, the trend of the "Arc" and "Spark" intensities, as given in the MIT tables, was taken into cognizance for selection of the "strong" lines.

The rare-earth elements have a second ionization potential around 11-12 eV, therefore at the temperatures of the Ap stars, particularly in the hotter Ap stars, most of the rare-earth elements are in the doubly ionized state. Swings (74) has identified lines of EuIII, GdIII, CeIII, SmIII, and LaIII in α^2CVn. Nearly all of these lines lie on the short wavelength side of $\lambda 3200\text{Å}$. But, the available set of 10 A/mm plates does not extend to the region below 3200Å . Strong lines of CeIII however, do lie in the

3200-3400 Å region, and therefore CeIII lines are included in Table V-1.

As we shall see in the following section and in the next chapter, lines of singly ionized rare-earths have not been observed in any of the Mn-stars, which usually have quite sharp lines. On the other hand, in many of the Si-4200 stars, which have B-V colors about as blue as the Mn-stars and have line width much greater, lines of EuII are clearly present. Therefore, an attempt was made to ascertain the presence of any rare-earths in the Mn-stars by examining the doubly ionized rare-earth lines (Swings 74) on the tracings of the 21 A/mm plates which do reach down to 3100 Å . However, at this dispersion, none of these lines is blend-free or sufficiently strong for a positive identification even in stars in which rare earths are known to be present.

Section V-B

Survey of the Rare-Earths

A search for all the rare-earths - except promethium, which does not have any stable isotopes - was made in seven Ap stars: α^2CVn, HD125248, γ Equ, HD188041, HD192913, HD153882, and π^1Boo. Strong EuII lines in the first four of these are well known; the presence of some other rare-earths in these stars has also been noted.

Therefore, these stars are very suitable for a survey of the complete list of the rare-earths.

HD192913 is listed as a "Si"-star in Bertaud's Catalogue, and Babcock (4) has noted the presence of EuII lines in it. The present study has revealed that this is a particularly interesting star because this is a Si-4200 star with strong lines of BeII. Furthermore, lines are fairly sharp and copious in number even though B-V = -.07. A large number of these lines can be identified with the rare-earths.

A search for the rare-earths was made in HD153882 (w = .4 Å) because the spectrum of this star is also rich in lines. However, it turned out that no noticeable line of any of the rare-earths is present (though this is listed as a "Eu-Cr" star in Bertaud's Catalogue). Amazingly, chromium is responsible for most of the profusion of lines in the spectrum of this star.

During the course of the survey of the Ap stars (Chapter III), lines of EuII were not observed in any of the Mn-stars even though most of the Mn-stars have fairly sharp lines. Therefore, the Mn-star π^1Boo was included in this program to find out whether some other rare-earth elements are present. In this star, no rare-earths have been found to be present.

Pattern of Distribution of the Rare-Earths. Estimates of the strength of all the lines in Table V-1 were recorded (for HD125248 only the ultraviolet plate was available). Then, the contribution due to some of the strongest rare-earth lines (marked, in order of preference by A and B in Table V-1) of an element was represented by an intensity parameter. (Such a value was chosen for this parameter that it was consistent with the strength of other weaker lines of that element.) It is to be emphasized that these intensities are not indicators of actual abundances because these depend upon the f-value of the observed line (differences in ionization potential, etc. produce effects of relatively lower degree). These intensity values tell us about the presence of the corresponding elements in a given star (In normal A-type stars none of these lines is observed at 10 A/mm.). In addition, these give some indication of differences in relative abundance of two elements between two stars. For example, from Figure V-1 we can easily infer that the $\frac{\mathrm{Gd}}{\mathrm{Tb}}$ ratio is much higher in HD188041 than in HD192913, or that the $\frac{\mathrm{Nd}}{\mathrm{Ce}}$ ratio is higher in HD192913 than in α^2CVn. Also, Ho is much stronger in HD192913 than in α^2CVn. Data on the ionization potentials of the rare-earth elements is seriously incomplete; in some cases even the first ionization potentials are unknown. However, values of the first, second, and probably the third ionization potentials are likely to be roughly the same for each of these rare-earths. Therefore, the above-mentioned marked differences in the ratios are probably not entirely due to differences in the ionization potentials and temperature.

Even though this study is of a qualitative nature, it clearly demonstrates the existence of differences between relative abundances of the various rare-earth elements in different Ap stars. A quantitative differential analysis of abundance of various rare-earths among the Ap stars has not been made; in their recent paper, Fowler et al (36) have not taken these differences into account.

TABLE V-1

Principal Unblended Lines of the Singly Ionized Rare-Earths in the Blue and Ultraviolet Regions of the Spectrum

λ	Intensity RMT	MIT Arc	MIT Spark	Neighboring Lines
LaII				
3265.67B	600	300	200	FeI 5.62; VII 5.89; CrII 6.25
3344.56	200	300	200	
3380.91	300	200	100	TiII 0.28; NiI 0.57; FeII 1.00
3612.34	50	8	15	ZrII 1.90; NiI 2.74
3949.10	600	1000	800	FeI 8.78; AI 8.98; PrII 9.44
3995.74	400	600	300	FeI 6.00
4123.23A	400	500	500	
4238.38	400	500	300	FeI 8.03; CrII 8.69; MnII 8.79
CeII				
3485.05	400	30	10	HoII 4.84 (See under HoII; FeI 5.34)
3534.05	300	35	10	
3560.80	500	300	2	VII 0.59; FeI 0.71
3577.46A	500	300	12	ZrII 6.88; MnI 7.88 CaII 8.03
3655.85	500	25	12	FeI 5.47; ZrII 5.56; GdII 6.15; Sm 6.22
3999.24B	500	80	20	ZrII 8.98; CrII 9.00, 9.07
4133.80	500	35	8	NdII 3.36
4186.60	600	80	25	CrI 6.36; ZrII 6.70; FeI 87.04
4628.16	500	20	20	CrI 8.47
CeIII				
3228.56	400			
3427.33	125	12	10h	
3443.61	150	8	15	
3459.37	200	2h	15wh	
3454.37	150	10	40	
3470.89	300	3	10	
3504.60	100	5	10wh	
PrII				
3851.62	200	200	150	GdII 0.97; NdII 1.75; SmII 1.88
3877.23	200	125	80	FeI 8.02
3908.63	150	100	50	FeI 7.94; CeII 8.41; CeII 8.54
3908.43	200	100	60	
3994.83	200	300	25	NdII 4.68
4118.48	200	250	50	CeII 8.14; FeI 8.55; SmII 8.55
4408.84	200	125	100	FeI 8.42
4496.43A	250	200	125	CrI 6.86; ZrII 6.96; MnII 6.99

Table V-1, Cont.

NdII				
4061.09	200	40	30	ScIII 1.3; CrII 1.77
4156.08	250	10	20	ZrII 6.24
4303.57A	400	100	40	FeII 3.17
4446.39	200	100	50	FeII 6.25
4451.57	400	100	50	FeII 1.55
4462.99	250	60	20	CrI 2.77
SmII				
3568.27	1500	40	50	ScII 7.70
3592.60	1200	40	50	VII 2.01; GdII 2.71
3609.49A	1200	60	100	NiI 9.31; CrI 9.48; CeII 9.69
3634.29	1500	100	25	CrII 4.04; FeI 4.33; FeI 4.70
3661.37	1000	100	50	VII 1.38; CrII 1.44
3670.84	1000	100	50	NiI 0.43; CrII 1.12; GdII 1.20 ZrII 1.28
3693.99	1200	100	150	FeI 4.01; YbII 4.19 (See 3694.19 under YbII)
3756.41	600	25	10	CrII 6.55; FeI 6.94
3885.29B	1000	50	50	CrI 5.08; CrI 5.22; ZrI 5.41; FeI 5.51
3896.98	600	50	50	CeII 6.80
3922.40	800	60	60	
4390.86	600	150	150	MgII 0.59; FeI 0.95
EuII				
3688.42	1500	1000	500	CrII 8.01; NiI .42
3724.94	4000	250	50	MnII 4.81; FeII 5.30
3819.64	6000	500	500	CrI 9.56; HeI 9.61, 76
3907.16	3000	1000	500	FeI 6.48
3930.50A	4000	1000	400	FeI 0.30; FeII .31; YII .66; CrII .88
3971.98	4000	1000		
4011.69	100	25		NdII 2.25; TiII .37
4017.58	100	25	25	FeI 7.16; CrII .96; MnI 8.10
4129.73	5000	150	50	CrI 9.21; GdII 0.37 (Listed under GdII)
4205.05	6000	200	50	CrII 4.66; CrII .83; VII 5.08; FeII .48
4435.58	3000	400	100	
GdII				
3362.23B	10000	150	180	ScII 1.94; YII 2.00; CrI 2.21; TmII 2.62; FeII 2.76
3439.99	6000	70	50	AlI 9.35; GdII 9.78; FeII 0.25
3481.29	5000	150	150	ZrII 1.14; CrI 1.30; CrI 1.54; GdII 1.80

Table V-1, Cont.

GdII, Cont.				
3894.70	2000	150	80	NdII 4.63; CeII 5.11; CrII 5.12; CrII 5.16; FeI 5.66
3916.51A	3000	150	100	ZrII 5.94; LaII 6.05; CrI 6.24;
3996.32	800	100	100	VII 6.42; FeI 6.73
4130.37	3000	200	10	EuII 9.73; CeII 0.71
4184.25	2000	150	150	SmII .76; LnII 4.26
4347.32	400	100	100	
4436.22	200	30	100	
TbII				
3509.17A		200	200	ZrI 9.32; CoI 9.84
3561.74		200	200	NiI 1.75
3776.49	(340)	100	100	TiII 6.06; FeI 6.45; YII 6.56
3848.76		100	200	MgII 8.24; NdII 8.52; CeII 8.60; SmII 8.78; CrI 8.98
3874.19B2		200	200	FeI 3.76; CoI 3.95; CrII 4.41, 4.76; CrI 4.57
DyII				
3407.79		150	9	NiII 7.30; FeI .46; GdII .61; ZrII 8.09
3523.98		15	20	GdII 4.20; CrII .54; NiI .54
3531.70A		100	100	EuII 1.15; MnI 1.85, 2.00, 2.12
3872.13	(600)	300	150	LaII 1.64; SmII 1.78; FeI 2.50
3944.70	(850)	300	150	AlI 4.01; NiI .13; FeI .75; FeI .89
4000.48B	(650)	400	300	NdII 0.49
HoII				
3398.98	(900)	40	60	FeII 8.36; FeI 9.34; CrII .54
3416.46	(600)	30	40	FeII 6.02; GdII .95;
3456.00B	(1800)	60	60	CrI 5.28, .60; FeII 6.00; TiII .39
3484.84	(700)	40	30	CeII (See 3485.05 under CeII)
3748.17A_1		60	40	YII 7.55; TiII 8.01; CeII .06; FeI .26; FeII .49
3861.68A_2	(300)	40	20	EuII 1.18; SmII 2.05; CrII .17
3891.02	( )	200	40	TmII 0.53; NdII .58; NdII .94; SmII 1.21
4045.44	(600)	200	80	ZrII 5.63; FeI .82
ErII				
3499.11	(650)	18	15	FeII 9.88
3616.58B	(300)	30	20	EuII 6.15; FeI .57
3692.64A_1	(700)	20	12	SmII 2.22
3830.53	(320)	8	1	SmII 0.29; PrII .72
3896.25A_2	(420)	30	15	FeII 6.11; CaII 6.80
3906.34	(850)	25	12	NdII 5.89; FeII 6.04; FeI .48; EuII 7.10

Table V-1, Cont.

TmII				
3241.53	200	150	125	ZrII 1.01; SmII .16; CrII .38; FeII .69; TiII .98
3362.62	300	250	200	GdII (See 3362.23 under GdII)
3425.08	300	200	300	GdII 4.59; EuII 5.02; CrII .09
3441.51	200	150	80	FeI 0.99; EuII 1.00; CeII .21; CrI .44; MnII .98
3462.20B	300	250	200	EuII 1.38; TiII .50; NiI .65; CrII 2.73
3700.26	300	150	80	VII 0.34; CrII .42; SmII .92
3795.76	600	250	150	FeI 5.00; GdII 6.37; ZrII .47
3848.02A	1000	400	250	SmII 7.5; NdII 8.23; MgII .24
YbI				
3987.98		1000	500	LaII 8.51
YbII				
3694.19	1000	500	1000	SmII (See 3693.99 under SmII)
LuII				
3397.07	150	50	20	EuII 6.58; FeI .98; TmII 7.50;
3554.43A	200	50	150	FeI 4.92; CeII .99 FeII 5.08
3976.65B	100	50	100	CII 6.19, .41, .67, 7.05
4184.26	120	100	200	GdII (See 4184.25 under GdII)

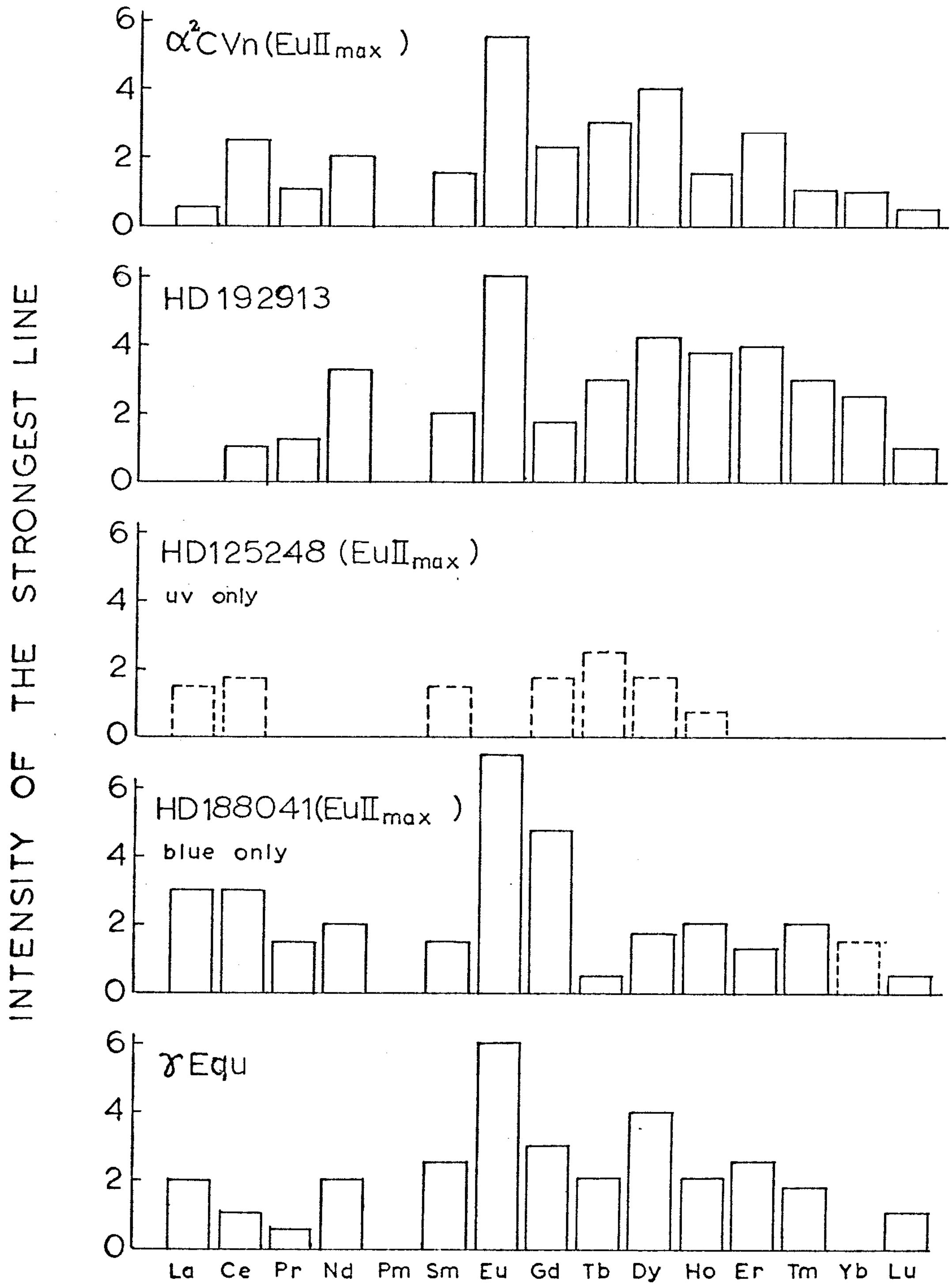
INTENSITY OF THE STRONGEST LINE
α^2CVn (EuII$_{max}$)
HD 192913
HD125248 (EuII$_{max}$)
uv only
HD188041(EuII$_{max}$)
blue only
γ Equ
0
2
4
6
La Ce Pr Nd Pm Sm Eu Gd Tb Dy Ho Er Tm Yb Lu

Figure IV-1

CHAPTER VI

DISCUSSION

An answer to "what makes the Ap stars different from normal stars" remains our ultimate goal. However, before a diagnosis can be made, symptoms must be well understood. Therefore, our first task is to discover systematics of behavior in all the groups of Ap stars. Patterns or peculiarities are quite diverse. Not only that the same peculiarities are not observed in the different types of Ap stars, often antithetical behavior is observed; e.g., silicon is very strong in some Ap stars, and weak in others. Such observations provide valuable constraints for explanations put forward.

In Part I of this chapter we shall present the observed systematics as discovered from spectroscopic work in the present study by discussing (a) the behavior of various elements in Ap stars, (b) correlations between observed anomalies, and (c) classification of Ap stars. In Part II some of the possible explanations will be discussed in light of these results.

Part I

BEHAVIOR OF THE VARIOUS ELEMENTS IN AP STARS

Behavior of the various elements among the different peculiarity groups is summarized in Table VI-1 . The following notation is used: N means about normal; + sign means strengthening, which if interpreted as abundance effect, represents an estimated logarithmic overabundance by 1/2 to 1; ++ sign denotes "overabundance" by large factors; - and -- are corresponding symbols denoting weakening or "underabundance"; a question mark denotes that either because of a lack of observable lines or because of the unavilability of suitable observational material, nothing could be said about those elements. Symbols given in parentheses are less certain and represent a suspected trend. When more than one symbol is used to describe an element, the first represents the more frequent behavior. Sargent (65) has given an analogous tabulation in Table VI (hereafter referred to as T6) of his review article. In our tabulation, some of the departures from T6 are indicated: an asterisk on an entry indicates that nothing was said about that in T6; major departures from T6 are underlined. Descriptions of the various elements are given below. Departures from results in T6 are also commented upon.

Be A detailed discussion of Be has already been presented in Chapter IV. Beryllium has been observed not only in some Mn-stars but also in some of the Si-4200 stars. This is a very important departure from the findings of Sargent et al (68). In T6, Be is described as "N,-" in contrast to our finding of "N,+" for the Si-4200 stars.

It is interesting to note that, so far, large overabundances of Be have been found only in hotter Ap stars.

Mg In the majority of Ap stars, Mg is not very much different from normal. But, in Si-4200 stars, as a group, Mg is on the weak side. However, regarding any possibility of a Mg-Si correlation, it is to be noted that cases of magnesium deficiency are found both among stars in which Si is strong (e.g., Si-4200 stars HD205087, HD175774, and HD224801) and among stars in which Si is weak (e.g., HD125248 (Cr-Eu) and HD8441 (Sr-Cr-Eu)).

In contrast to T6, weakening of Mg in some Cr stars is observed.

Si Silicon is very strong in Si-stars, normal or moderately strong in Mn-stars, and mostly weak in Sr-stars. The Cr-group includes stars both with Si strong, and weak.

In T6, Si is listed as mostly normal for Sr stars. As discussed by Searle and Sargent (69), equivalent widths of the SiII lines vary only slowly along the spectral range of the Ap stars. For some of the Sr-stars only low dispersion plates were available for the blue region, in which SiII lines are observed. However, Si seems to be consistently weak in these Sr-stars. Further details will be discussed later.

P Bidelman (14, 65) has identified PII lines in the Mn-stars κ Cnc and 112 Her. In no other Mn-stars of this survey, PII lines have been observed. In μ Lep, PII lines if present, are very weak.

Ca Lines of CaII are weak or very weak for their temperature in all the "Cr" and many "Si" and "Sr" stars. (A negative correlation between Cr and Ca is discussed later.)

In T6, Ca is shown as "-" for Mn-stars. However, as illustrated by the K-line in "Mn" and normal star, in Figure VI -3 , Ca is fairly normal in these Mn-stars.

A very unusual K-line profile, with very wide and shallow wings, is observed in HD188041 and γ Equ, and to a lesser degree in 52 Her. These and peculiar profiles of K-line in other Ap stars have been mentioned in the literature (e.g. 59, 5).

<u>Sc</u> Considerable weakening of ScII lines has been found to be a general characteristic of Am stars (Preston 63 , Conti 23). However, in most of the Ap stars of the present survey which are not too hot and have lines not too wide for detection, Sc (as inferred from ScII lines) is strong. Even in several hotter Ap stars, ScII lines are present indicating a considerable strengthening of Sc (in normal stars with the same B-V colors, these are not present).

In γ Equ, Sc shows weakening relative to Fe as indicated by the ScII 4246.83: FeI 4247.43 ratio. In some other stars in Table III-4(a-d) , Sc is indicated as possibly weak; but for these stars, high dispersion spectrograms were not available.

<u>Ti</u> The ratios TiII 4171.90: FeII 4173.43 and TiII 3685: CrII 3677 show a great strengthening of Ti relative to Fe and Cr respectively in Mn-stars (see Cr-Ti correlation). Many lines of Ti II which are absent or very weak in normal stars in the B-V range of the Mn-stars are present with fair strength in Mn-stars. It is very safe to conclude that Ti is very strong ("overabundant") in Mn-stars;

in T6 it is listed as normal.

In other peculiarity groups, in the majority of stars, Ti does not depart markedly from normality. However, there are small deviations in several of these stars. Some Cr-stars, e.g., HD153882 and HD9996 (and possibly the Si-4200 star HD205087), show very conspicuous weakening of TiII lines. As an illustration, attention is drawn to TiII 4463.76 and FeII 4555.89 lines in Figure VI-1 : HD153882 and HD9996 show weakening of Ti very clearly. In these other groups, no examples of any outstanding strengthening of Ti lines have been found.

Cr All the groups, except the Mn-group, show great strengthening of Cr lines. We have already shown that the known physical processes such as micro-turbulence, magnetic intensification, suppression of the double ionization, etc., cannot account for this strengthening. If interpreted as an abundance effect, these would yield overabundance factors of possibly between 5 and 50 or more. In Mn-stars, Cr is about normal, or in many cases weak as indicated by the equivalent width measurements of CrII 3403.22. In these respects, our results differ from those listed in T6.

<u>Mn</u> In the Mn-stars, MnII lines are very strong. Also in many Cr-stars MnII lines are quite strong (See Mn-Cr correlation). In other Ap stars, MnII lines are either normal or somewhat strong.

It is to be noted that in some Si-4200 stars (e.g., 4 Cyg), lines of SiII, CrII, and FeII are greatly strengthened, but not those of MnII. This is further confirmation that micro-turbulence or magnetic intensification cannot be a major cause of the strengthening of lines in these stars.

<u>Fe</u> For relative abundance determinations, usually, Fe is used as comparison. In the present work, however, an attempt has been made to make estimates relative to hydrogen (see Section III-D). In Mn-stars, Fe is generally on the weak side, as discussed under the Mn-Fe correlation. In other peculiarity groups, Fe is mostly normal or strong. Strengthening of FeII 4555.89 relative to TiII 4563.76 , in several Cr-stars, is to be noticed in Figure VI-1 .

<u>Ni</u> Using πCet (B7V) , θAql (B9.5III), αCMa(A1V), θPeg (A1V), and θLeo(A2V) for reference, nickle is found to be weak in most of the Ap stars. In Mn-stars lines are sharp, and for most of the Mn-stars of this survey, high dispersion spectrograms are available; all these show weakening of NiII lines relative to the

normal stars having corresponding B-V colors. Some of the Si-4200 stars have quite wide lines, but in other stars of this group, NiII lines are on the weak side for their B-V colors. Of the "Cr" and "Sr" stars for which high dispersion plates are available, only in γ Equ, Ni may be on the strong side; in others it is weak or normal.

It is to be noted that $\left[\frac{Ni}{Fe}\right]_{*} - \left[\frac{Ni}{Fe}\right]_{\odot} = +1.0$ for α Lyr (Hunger 41), and +.3 for α CMa (Boyarchuk 19). Furthermore, inspection of spectrograms indicates that Ni abundance in the sharp-line star Peg may even be higher. Therefore, the question "whether the sun or the above listed normal stars represent normal Ni abundance" deserves further investigation.

<u>Ga</u> Bidelman has identified lines of GaII in 3Cen A, and in the Mn-stars χ Cnc and 112 Her. In the present survey, GaII lines have been examined in six Mn-stars for which blue-plates were available. Of these, three (χ Cnc, π^1Boo, and μ Lep) have strong GaII lines; in υ Her GaII lines are perhaps present, but quite weak; in the other two, ι CrB and ϕHer , GaII lines are not noticeable (In normal stars GaII lines are not noticeable.).

<u>Sr</u> Strontium is always either strong or normal in all the types of Ap stars. Besides the Sr-stars, all the Si-stars, several "Cr" stars, and some Mn-stars show

strengthening of strontium. In the rest of the Ap stars it is normal or strengthened by fmall factors. Strengthening of SrII lines in Si-4200 stars is illustrated in Figure VI-2 .

<u>Y</u> In spite of their high temperatures, many lines of YII are clearly present in the blue as well as UV regions in the spectra of the Mn-stars (only in κ Cnc, which is an atypical Mn-star according to Sargent and Searle (67), yttrium is not strong enough to show any YII lines). In Figure VI-3 , YII $\lambda\lambda$3950.4 and 3982.6 are clearly present in the Mn-stars while in α Lyr they are not visible at all and in α CMa only barely visible; α Lyr and α CMa are much cooler than the Mn-stars. From these observations, great strengthening (i.e. "overabundance") of yttrium in the Mn-stars is quite certain (See Mn-Y correlation.).

In other groups, it is moderately strong or normal; in two "Cr" and one "Sr" stars it is weak. Though in most of the Ap stars both Sr and Y are on the strong side, there is no evidence for a strong, positive correlation between the strengthening of the two elements. It is interesting to note that in none of the Sr-stars is Y greatly strengthened - it is either normal or only moderately strong.

<u>Zr</u> Zirconium is strong or normal in most of the stars of all the peculiarity groups. In two Cr-stars,

it is weak however (See notes on observability of Zr in Section III-C).

Ba See Ca-Ba correlation.

The Rare-Earths Europium is strong in many of the Si-4200 stars, but in none of the Mn-stars. (Mn-stars lie in the same B-V range as the Si-4200 stars, and have much sharper lines.). The entry (N,+) in T6 for Mn-stars is therefore not substantiated by the present observations. In most, but not all, of the "Cr" and "Sr" stars, Eu is overabundant by large factors; a conspicuous exception is HD153882. Other details on the rare-earths have been discussed in Chapter V.

CORRELATIONS BETWEEN OBSERVED ANOMALIES

Correlations between anomalous behavior of elements are investigated, in this first study, for a large number of elements. Correlations between a few selected elements and peculiarity types have been studied earlier (Sargent, Searle 67, 69).

To discuss correlation, the "abundance defects" (as given in Tables III-4 (a-d)) were plotted for most of the possible combinations of those elements. These plots revealed some very strong and other less-strong correlations, which are discussed here:

(a) Mn-Y A very remarkable result is that in all the Mn-stars (except χ Cnc) of this program yttrium is very strong. Figure VI. shows this correlation. Figure IV-6 illustrates strengthening of YII 3600.7 and Figure VI-3 illustrates strengthening of YII 3930.4, 3982.62 in Mn-stars. (In normal stars, lying in the same B-V range as the Mn-stars, lines of YII are not noticeable and, therefore, are on the linear part of the curve of growth.)

In the B8 star HD22401, Kraft (47) has found the presence of YII lines, which are not present in normal B8 stars. To see whether Mn is strong in this star, on the blue plate (No. Pd8432) of this "yttrium star", which was kindly lent to me by Dr. Kraft, I have examined the MnII lines which are found only in Mn-stars. Indeed, it seems likely that these MnII lines might be present and that this is an Mn-star. However, additional spectrograms, preferably in the UV, are needed for a confirmation.

In the other peculiarity groups, neither Mn nor Y departs from normal by large factors.

(b) <u>Ca-Ba</u>: In all those cases in which it was possible to make estimates of BaII λ4554, a very systematic correlation between these lines and CaIIK was observed as can be seen from Figure VI-10 which is based on the results of more than 20 stars representing all the peculiarity groups. (Blue plates were not available for the rest of the stars of this program, and therefore the strength of BaII λ4554 could not be estimated for them.) This correlation is very well illustrated in Figure VI-1 in which normal and Ap stars are arranged in a B-V sequence. All the stars in which CaIIK is weak, also show weakening or absence of BaII 4554. In the Mn-star ι CrB, CaIIK is normal, and BaII 4554 can be seen, which means that Ba is not weak. (HD188041 has wide shallow wings of K-line, and it is difficult to estimate the line strength from high dispersion plates. Therefore, for this star, K-line strength was estimated from a 40 A/mm plate.)

(c) Cr-Ca: A fairly strong negative correlation exists between Cr and Ca - which means one is strong and the other is weak - for stars of all the groups. In "Si", "Cr", and "Sr" groups, Cr is strong and Ca is mostly weak; in Mn-stars Ca is normal (or possibly strong), while Cr is normal (or probably weak. This correlation is shown in Figure VI -11 . In Figure VI-1 we notice that all the stars with CaIIK weak, have strengthened CrII lines; many lines of CrI and CrII absent in normal stars are present in the spectrum of these stars.

(d) Ca-Sr: The existence of a strong correlation between Ca and Ba, which lie in the same column of the periodic table, is discussed above. Strontium falls in the same column and is intermediate between Ca and Ba. Therefore, any possibility of a Ca-Sr correlation will be of particular interest.

Figure VI- 12 does not present a *prima facie* case for the existence of a Ca-Sr correlation. But, in this figure, if we disregard the Mn-stars (for which, as discussed later, there is strong evidence that these are greatly different from all other types of Ap stars) we notice that the scatter of the rest of the points suggests the possibility of a line with positive slope. This line is shifted upwards and does not pass through the origin.

Only as a possibility, it might be conjectured that this behavior of Sr is caused by two mechanisms: One which increases strength of the SrII lines because of a real overabundance of this element in the observable part of the stellar atmosphere. The other mechanism effects Ca, Sr, and Ba in an analogous manner reducing the observed strength of lines of (singly ionized state of) these elements thereby manifesting a positive correlation between them. The possibility of a physical mechanism for producing such a correlation will be discussed later.

It is to be noted that, in HD8441, in which both Cr and Sr are very strong, the negative Cr-Ca relation prevails, and CaIIK lines are very weak.

(e) Cr-Ti: In Ap stars, ratios of the strength of TiII and CrII lines show considerable departure from that in normal stars. Comparison of TiII 3685 and CrII 3677 in Figure VI-8 ; and TiII 3383 and CrII 3382 in Figure VI-7 very clearly demonstrates these departures relative to normal stars of the same color. In Mn-stars the Ti/Cr ratio is larger; in the other groups, (in which Cr is strong), this ratio is smaller than in normal stars. These departures suggest a negative Cr-Ti correlation of the following nature: In many stars in which Cr is strong, Ti is weak; in Mn-stars, in which Ti is very strong, Cr is normal or weak. This correlation is, probably, not as strong as some others discussed earlier.

(f) Mn-Ti: Figure VI-13 suggests a positive correlation between Mn and Ti for Mn-stars. This correlation is quite secure; both Mn and Ti are strong.

(g) Mn-Cr In Figure VI-14 we notice that Mn is strong in two types of stars: One, in which Cr is normal or weak (Mn-stars); and the other in which Mn is strong to a lesser degree, but Cr is very strong. Therefore, firstly, it is to be recognized that all stars with Mn strong are not necessarily members of the Mn-group. Secondly, for stars of the Mn-group, in which Cr is on the weak side, the suggestion by the Burbidges (21) that Mn and Cr are overproduced from Fe, is not applicable.

(h) Mn-Fe: Figures VI-4 and VI-14 illustrate the weakening of FeII lines in Mn-stars. Also, a comparison of FeII 4555.89 and CrII 4558.66 (FigureVI-4) shows that in Mn-stars Fe is weaker relative to Cr; as already discussed. Cr is not strong in these stars

Spectrum Variations

The relationship between the behavior of various elements is not always unique; sometimes relationships are even opposite in different stars. Such diversities make the task of finding the explanation of the phenomena of Ap stars and spectrum variables more difficult. However, these diversities must be taken into account in any satisfactory model. Following are some examples of such diverse relationships.

(i)Diversity of relationship between the behavior of SiII and CrII lines in HD125248, α^2CVn, and HD153882, which are alpha-variables of Babcock's magnetic variable classification: In HD125248, SiII lines are very weak at EuII max. phase, and only somewhat weaker than normal at CrII max. phase. In α^2CVn, SiII lines are quite strong at all phases though Si and Cr vary in the same phase, as is the case with HD125248.

In HD153882, variations in only SrII lines have been reported in the literature. The three 10 A/mm plates of this star, taken for the present work, indicate that SiII lines are variable; while CrII lines are constant and always strong, SiII lines undergo variations from very weak to normal. (Figure VI-5)

(ii) Diversity of correlation between behavior of various elements is further demonstrated by these examples: Deutsch (26) found that in ι Cas, CaII and CrII lines undergo variations with the same phase, but in χ Ser, CaII and CrII lines vary in opposite phase (Deutsch 28). In α^2CVn, HD125248, etc. EuII and CrII vary in opposite phase, but in 73 Dra, according to Farraggiana and Hack (34), EuII and CrII vary with the same phase. Furthermore, as was found by Babcock, in HD125248 magnetic field maximum and EuII max. occur in phase, but in α^2CVn these are at opposite phases.

CLASSIFICATION OF THE AP STARS

Old Classification

For presentation of results in Tables III-4 a-d , a system of four peculiarity groups, "Mn", "Si", "Cr", and "Sr", has been used; this is a modification of the earlier systems. Reasoning for these modifications lies in the characteristics discovered in the present study. A brief description of these four groups is as follows:

(a) "Mn": Some of the Mn-stars were earlier, in the HD Catalogue or other low dispersion work, classified as "4128-31" or "Si" type. The reasons for this are that in Mn-stars, also the SiII 4128-31 lines are strong and that Mn-II lines are very weak in the blue region of the spectrum.

Of the Mn-stars in Table III-4a all, except 87 Psc, are known as of Mn-type in the more recent literature. Star 87 Psc is listed as "Si" type in Bertaud's Catalogue. However, its UV spectrum shows all the typical Mn-star characteristics, (a blue plate was not available for this star). Consequently, this star is listed under the Mn-group.

(b) "Si": Stars usually classified as Si-4200, "Si", and Si-Sr and in which Si is genuinely strong* are all combined into this group. This includes α^2 CVn which is usually designated by Eu-Cr. The star γ Equ has Si strong, therefore it also should be included among the Si-stars. However, for the present, it is retained among the "Cr-stars".

(c) "Cr": This group includes stars in which chromium shows most conspicuous strengthening.

*In the present study, unless mentioned otherwise, the term "Si-star" has been used only for those stars in which SiII lines are too strong for their B-V color and in which their strengthening is not caused by a low value of the electron pressure.

(d) "Sr": Stars in which strontium lines show most conspicuous strengthening are put under this group.

New Classification

The above system used for presenting results in Table III-4(a-d) is only an amended version of the usual system. It is possible that some other way of classification may be more useful and may contain a clue to a better understanding of these objects. Before proceeding to suggest some alternative systems, attention is drawn to the following facts:

(i) Initially, "abundance" results (Tables III-4) for all the stars were arranged in a B-V color sequence irrespective of their peculiarity type. But no systematic pattern emerged from this arrangement. This strongly suggests that no strong temperature-peculiarity type relation exists which encompasses all the peculiarity types. Consequently, the temperature-peculiarity type relation in the form found by Morgan (57) and the Jascheks (42) is fortuitous, and it does not reflect existence of a physical cause in the Ap stars which, for different temperatures or spectral types, produces different peculiarity types.

(ii) On the other hand, it was noticed that, unlike most of the other cooler Ap stars, γ Equ has strong Si lines. This, and other features in the spectrum of γ Equ suggest that, chemically, this star might belong to the same general group as the other hotter Si-stars. Sargent (66) has expressed the view that α^2CVn, β CrB, and γ Equ (B-V = -.11, +.27, and +.27 respectively) have probably quite similar abundances.

If in the future, more detailed investigations affirm the finding that γ Equ indeed belongs to the Si-group, its implication would be, that, instead of arranging groups of Ap stars sequentially in a single temperature sequence (as was done in the above mentioned works of Morgan and the Jasckecks), we should split the Ap stars into more than one parallel sequence which may or may not extend over the entire temperature range of the Ap stars.

(iii) A perusal of the characteristics of the Ap stars leads us, immediately, to separate the Mn-stars from the rest. Some of these striking differences are that in Mn stars CaII-K is normal, rare-earths are not strong, and Cr is on the weak side. For other types of Ap stars, mostly the opposite is true. Other differences such as agreement of (B-V) vs. θ_{ion} relation with normal stars is discussed in Chapter II. Further, though the B-V range of Mn-stars lies within that of the Si-4200 stars, the two groups are spectroscopically very different from each other.

(iv) Though most of the Si-stars are hotter Ap stars, we have noticed the extension of this group to the cooler region (γ Equ has B-V = +.27); the deblanketed value of B-V is probably between +.1 and +.15). For a search of an analogous extension of the Mn-group, cooler Ap stars (B-V > 0) were checked. But, none of the cooler stars of this survey was found to share the characteristics of Mn-stars. Thus, it seems that Mn-stars belong to quite a localized group with B-V around -.10 and with weaker examples down to B-V -.05.

(v) After eliminating the Mn-stars, the remainder of the stars - "Si", "Cr", and "Sr" groups - share some common characteristics, e.g., Cr is strong in all these; rare-earths are present in most (though not all) of them; except for some "Sr" stars nearly all have very weak K-line. On the other hand, there are some very striking, dividing characteristics among stars of these groups (now we are referring to Ap stars other than those of Mn-type): Silicon is <u>strong</u> in the hotter stars, which are "Si"-stars, but <u>weak</u> in most of the cooler stars, which include all the stars of "Sr" group and most of the "Cr"-stars. Searle and Sargent (69) have shown that the equivalent width of λ4128 and λ4131 of SiII vary only slowly over the B-V range of Ap stars. Consequently, the above-mentioned very conspicuous strengthening, and weakening of SiII lines cannot be explained by an uncertainty in the temperature of these stars. Spectrograms of stars with SiII lines strong, weak, and normal are shown in Figure VI-5 . In this figure, both $\lambda\lambda$4128 and 4131 lines of SiII are prominent in the Si-stars HD34452, 108 Aqr, and 21 Per and in the normal stars ι Her, πCet, oPeg, αCMa, and θLeo. Also in γ Equ, a cooler Ap star, SiII lines are strong. But in other stars λ4128 is blended, and only λ4131 is reliable. Weakening of this line in HD188041, 52 Her, HD9996, and HD8441 is very conspicuous. In HD125248 and HD153882, SiII 4131 is weak on certain phases.

In view of the above discussed facts and characteristics of the various groups, a new scheme of classification is proposed here. The basic difference between this scheme and the earlier classification-schemes is that in the present stars are classified according to a general pattern of behavior of several elements in that star. In the earlier schemes, on the other hand, each feature or each element defined a different peculiarity type; the consequence of this was that stars had to be divided into a large number of categories (Osawa (62) has classified Ap stars into 13 categories!).

Sequence A: CaII-K is not weak (normal or strong); Mn, Y, and Ti very strong; Si on the strong side; Cr and Fe on the weak side (or normal); the rare-earths (specifically Eu) not present; Sr strong or normal. Sequence A comprises the Mn-stars of the earlier classification.

Sequence B: Moderate to great strengthening of Cr and Sr; CaII-K is often very weak for its B-V color (particularly in stars with Cr very strong); in most cases rare-earth elements are present. The "Si", "Cr", and "Sr" groups of the earlier classification comprise sequence B.

Stars of sequence B may further be sub-divided on the basis of the strengthening or weakening of Si:

B1: Silicon <u>strong</u> stars: All the Si-stars and stars like γ Equ, which are classified under "Cr" stars but in which Si is strong, fall in this category.

B2: Silicon <u>weak</u> stars: Probably all the "Sr"-stars and most of the "Cr"-stars have SiII lines too weak for their temperature inferred from their B-V color. Such stars fall in this group.

It is to be noticed that the Si-stars, which on the average are the hottest, fall in sequence B1; Sr-stars, which on the average are coolest, fall in sequence B2; and

Cr-stars, which are intermediate, belong to B1 or B2. This makes one suspect that the differences in behavior of Si in the stars of sequence B are due to an effect of temperature. However, this does not seem to be the case because we do not observe a systematic trend according to which hotter Cr-stars (therefore closer to Si-4200 stars in the B-V sequence) show strengthening of Si, and the cooler ones (closer to Sr-stars in the B-V sequence) show weakening of Si.

The above classification system is schematically shown in Figure VI-16

Besides the strength of Si, there are some other prominent characteristics which might be used for subdividing the stars of sequence B. Though in most of these Ap stars which are not too hot and not too wide-lined, strong EuII lines are observed, there are some e.g., HD153882 and possibly HD9996, which are important for the absence of EuII lines. Thus the presence of strong rare-earth lines may provide another criterion for a subdivision. However, the number of such stars with EuII lines absent is very small. Another prominent feature is the weakening of the CaII-K line in a high proportion of these stars; in many Sr-stars and a few Si-4200 stars it is not weak. This may provide still another criterion for the subdivision of sequence B.

Of the above criteria, the one based on the strength of Si is perhaps most profound because among Si, Eu, and Ca, silicon is the most abundant element.

Part II

SOME POSSIBLE PHYSICAL CAUSES OF LINE STRENGTH ANOMALIES

We have already discussed the fact that, for many of the observed anomalies of line strengths, effects of temperature, electron pressure, opacity, micro-turbulence and magnetic intensification provide insufficient explanations. The diversity of the observed features is such that different features require different combinations of these parameters; and, except for a few marginal examples of Ap stars, it is not usually possible to find a single combination of these parameters which would explain all the peculiarities. We have discussed the presence of YII lines in Mn-stars, or the abnormality of the CrII:FeII ratio in many Cr-stars as obvious examples of such difficulties. Furthermore, as is shown by Struve and Swings (72) and subsequently discussed by the Burbidges (20), it is not possible to explain spectrum variations in α^2CVn by any appropriate variation in the physical parameters. The observed variations in micro-turbulence (if real) and the variations in temperature as inferred by Provin would produce changes only in the wrong direction to explain spectrum variations in α^2CVn(20).

Some authors have attempted to explain the observed peculiarities by invoking other types of physical effects. Here, we shall compare our observations with their predictions.

(i) Effect of the magnetic field on the structure of the stellar atmosphere: -

It is assumed that the presence of magnetic fields will stop convective energy transport in the atmosphere of Ap stars altering the temperature gradients, and consequently altering the strength of observed lines. De Loore et al (25) have made such calculations by assuming a totally radiative atmosphere. They find, for stars with mass $M \lesssim 1.75\ M_{\odot}$: that (a) these lie about one magnitude above the normal main sequence on a temperature-luminosity diagram, (b) temperature-H_{γ} equivalent-width relation remains unaltered, and (c) the equivalent width of metallic lines (SiII, MgI, MgII, SrII, and CaII) are altered.

Firstly, it is to be noticed that these calculations are applicable to stars with $M \lesssim 1.75\ M_{\odot}$ (i.e. later than A5), while most of the Ap stars have earlier spectral types. Secondly, their calculations show an increase in strength of SiII 4131, CaII H and K, and slight weakening of SrII 4077 and 4125. These are just the opposite of what we observe in the cooler

Ap stars. It is to be further noted that according to Mihalas (55), for normal stars, convective flux becomes important for $\theta_e \gtrsim .6$ (therefore, B-V $\gtrsim$ +.1). For $\theta_e \lesssim .5$ (B-V < 0), convective flux is only minor and does not produce appreciable difference in the temperature gradient, on which equivalent width of lines depend. Therefore, we can conclude that the stoppage of convection by magnetic field cannot produce a major change in line strengths in Ap stars with $\theta_e \leqslant .5$.

(ii) Non-LTE effects:-

Struve and Swings (72) proposed a mechanism for the weakening of CaII lines in Ap stars. They observed six abnormal A stars (Am) in the Hyades, and found, "in all the six stars, line CaI 4226 is present and is of an intensity similar to that observed in normal stars of spectral type A8 to FO." Earlier, Titus and Morgan (75), on the basis of CaII-K line, had assigned types between A1 and A5 for these stars. From this, Struve and Swings suggested, "it is probable that the weakening of CaII is not accompanied by a similar weakening of CaI, and that the effect cannot be caused by low abundance of calcium." Next, they proposed that, in these stars, it is possible that the reversing layer is surrounded by a chromosphere, L_β is in emission, and that the weakening of CaII-K is due to overionization of Ca^+(I.P. = 11.9 eV) by L_β radiation (12.05 eV).

In the sun, chromospheric activity and magnetic field are intimately correlated. With the discovery of strong magnetic fields in the Ap stars, the possibility of a chromosphere seems more plausible now than it was first suggested by Struve and Swings.

At the temperatures of the hotter Ap stars, lines of neither CaI nor of CaIII are observable, and we cannot establish whether weakening of CaII lines is due to underabundance of calcium or due to overionization of Ca^+. It is very interesting to note, however, that there is a similar coincidence between the ionization potential of Ba^+ (10.0 eV) and $L\alpha$ (10.15 eV), and that BaII lines are found to be weak in exactly those stars in which CaII lines are weak. This Ba-Ca correlation has been discussed earlier. This strong correlation suggests the possibility that weakening of CaII and BaII lines is produced by a common cause, possibly overionization of Ca^+ and Ba^+. (Both Ca and Ba are in the same column of the Periodic Table.)

But, even without making quantitative calculations, this explanation presents some difficulties. Lines of SrII are strong, to a varying degree, in almost all of the Ap stars, including those in which CaII and BaII lines are very weak. Since the continuous absorption coefficient varies only slowly (proportional to ν^{-3} for hydrogenic atoms), it would be expected that L_β radiation (12.05 eV) overionizes, though to a lesser degree, also Sr^+(.P. = 11.0eV) making SrII lines weaker:

We consider two hypotheses: (a) Ca^+, Ba^+, and Sr^+ are overionized by the physical process mentioned above. But Sr is intrinsically overabundant to an extent that the effect of overionization of Sr^+ is more than compensated. In this hypothesis, the intrinsic abundance anomaly of only one element needs to be explained, by process of nucleosynthesis or some other mechanism. (b) In the other hypothesis, it is assumed that the peculiarities of Ca, Ba, and Sr are all due to intrinsic abundance anomalies. In this second hypothesis we need to explain abnormal abundances of all the three elements with the additional constraints that Ca-Ba correlation is maintained.

Though, of the above two alternatives, the former seems to be simpler, one may ask, "is there any evidence that L_β is in emission?" There is evidence (18, 70) that in emission line stars, in which L_α and L_β are presumably in emission, equivalent width ratio of OI 8446:7774 departs (decreases) from the usual value of .6 because of a decrease in absorption in the λ8446 line. Sargent and Searle (67) find that OI lines are very weak in the cooler Ap stars. But, they state, "In each case the oxygen blends at λ7772-7775 and λ8446 behave similarly." However, they have given only the upper limits to equivalent width-values of OI lines; and, therefore, it is not clear whether along with an underabundance of oxygen the effects of $L\beta$ emission are present in these Ap stars. Nevertheless, it is interesting to note that nearly all the stars in which Sargent and Searle find weakening of OI lines show weakening of CaII-K also.

SOME OTHER EXPLANATIONS

From the findings of other investigators (72,69) as well as from the foregoing discussion, it is quite clear that, at least some of the abnormalities of line strengths are due to intrinsic abundance differences. Moreover, for line-strength variations in periodic spectrum variables, there is, so far, no adequate explanation based on variations in physical conditions only. Therefore, it would be useful to discuss, in light of the results of the present study, some of the proposals that require an uneven distribution of various elements over the surface of an Ap star.

(i) Aspect Effect

We have shown earlier that in Ap stars of all the peculiarity groups SrII lines are strong. In the sharp line stars, such as HD8441 (ω = .08Å), as well as in the broad line stars, such as 21 Com ($\omega = 1\text{Å}$), Sr is strengthened by large factors. If we assume that the observed Ap stars encompass a nearly complete range of inclination of the rotation-axis and that Sr-rich patches are formed by migration of Sr from other areas, we should expect Ap stars with Sr weak, as well as those with Sr strong. Consequently, if the Sr-rich patches exist, these must be due to migration of Sr from the <u>deeper</u> layers, or due to an over-production of this element on the stellar surface.

(ii) <u>"Optical Pumping"</u>

For the separation of elements, Babcock (6) has proposed a mechanism in which atoms experience a force $\mu \nabla \vec{H}$, μ being the magnetic moment of the atom and $\nabla \vec{H}$, the gradient of the magnetic field. Thermal collisions are a major disaligning cause; however, the flip resistance is particularly marked for atoms in states with spherical symmetry (S-states). He has calculated magnetic moments of atoms and ions in the S-ground state, and found that on this basis, besides some elements with very low abundances, Eu, Mn, and Cr will be concentrated most effectively .

Firstly, only the vertical gradients, not the horizontal gradients, can be effective in increasing the concentration of elements because horizontal movements can increase concentration on the observed hemisphere of the stars by a factor of two or so only. Large concentrations can be produced only if those elements are pumped from a deeper and thicker layer to a thinner layer on the surface.

Secondly, though the above-mentioned mechanism may be qualitatively consistent with the strengthening of Cr, Eu, and Mn in the case of some of the Cr-stars such as HD2453 and 17 Com A, it does not explain the strengthening of Sr. Observationally, in Mn-stars, Eu is not strong and Cr is possibly weak, while YII lines are very strong (the ground state of a YI is $^2D_{3/2}$ and of YII ^1S_o). These abnormalities cannot be explained by this mechanism. Also, the weakening of O in many Ap stars (67), strengthening of Si in Si-stars, weakening of Si in many "Cr" and "Sr" stars, strengthening of Be in several stars, etc., cannot be explained by this process.

(iii) Accretion of Matter from Binary Companion

To explain the underabundance of Be in α CMa, Bonsack (17) suggested that α CMa is covered by a layer of Be-deficient material ejected by its white dwarf companion during the course of its evolution. In α CMa, metals are about ten times overabundant, relative to hydrogen, than in α Lyr. If one extends this possibility to explain the Ap stars by assuming that these peculiarities reflect the composition of matter ejected from an undetected companion, the problem, in essence, is changed to that of finding ways in which nuclear reactions would produce the observed abundances. Moreover, then, it would not be possible to explain the overabundance of Be we have found in several of the Ap stars.

(iv) Separation of Elements

In 1932, Miss Farnsworth published the period of EuII variations in α^2CVn using observations of the preceding 17 years. In 1947, Deutsch found that his observations were in agreement with this period. Fifteen years later, again, the phase of my plates agreed with that predicted by Miss Farnsworth's value of period. This remarkable constancy of the period of spectrum variations, over a few thousand cycles, suggests that the separation of elements is effective and enduring in the spectrum variables.

Therefore, after the failure of the hypothesis for explaining spectrum variations by periodic changes in physical conditions, the separation of elements is, now, a very important hypothesis, irrespective of the question of whether the abnormal abundances are produced by nuclear reactions or by the migration of elements.

NUCLEOSYNTHESIS PROCESSES

Because of the failure of all other explanations proposed so far, in which unusual physical conditions are invoked, it becomes necessary that other mechanisms, such as nuclear reactions in the outer layers of stars, should be investigated. However, the validity of a certain proposed nuclear process is to be judged from the agreement of its predictions with the observed facts. The correlations and other characteristics discussed earlier in this chapter provide such observed facts and constraints for future work on nucleosynthesis in Ap stars. Here we shall see whether the nuclear processes that have already been proposed agree with our observations.

(i) The observed large overabundance of Be in several Ap stars is perhaps the strongest evidence in favor of the occurence of nuclear reactions in the outer layers of Ap stars. Beryllium and other rare-light elements are quickly destroyed by thermal collisions with protons at temperatures higher than a few million degrees. Consequently, the cause of overabundance must lie in the upper ($T < 4 \times 10^{6\circ}$ k) layers of the stars. The low magnetic moment of Be (Babcock (6)) does not support the possibility that its observed concentration is caused by magnetic separation. Unless some other feasible mechanism of separation is proposed, nucleosynthesis is most likely to provide the explanation for Be overabundance. However, as yet no satisfactory process for the production of rare-light elements is known; even in the case of the sun the observed ratio of Be:Li presents difficulties to the proposed mechanism (Bashkin and Peaslee (9)).

(ii) In a recent paper, Fowler et al (35) have discussed some processes for the production of peculiar abundances in Ap stars. On the basis of an earlier paper they concluded:

(a) The large abundance excesses of the rare-earths demand a process of neutron addition; and

(b) The process cannot be entirely slow; otherwise Ba would be in excess, which is not the case, and Eu, Gd, Dy, Ho would not be overabundant.

In this context, it is interesting to note that in the cooler Mn-stars - e.g., ϕ Her and ι CrB - in which there is any liklihood of observing BaII lines, BaII 4554 <u>is</u> present and hence Ba is not weak (may be strong). Furthermore, in these Mn-stars there is no indication of an overabundance of Eu or other rare-earths. On the other hand, in stars with strong EuII lines, BaII lines are mostly weak (though there are also stars, e.g., HD153882 and HD9996, in which BaII lines are very weak <u>but</u> Eu lines are not present).

The following observations might be interpreted as an indication that, for formation of the rare-earths, the s-process was relatively more important in HD192913 than in HD18804: Both HD192913 and HD188041 have strong Eu (r- and s-process), but in the former, Gd(r) is only moderately strong, Nd(s) is quite strong, and Ba(s) is on the strong side; while in HD188041 Gd(r) is very strong, Ba(s) is quite weak, and Nd(s) is only moderately strong.

(iii) Fowler et al (36) have also discussed the case in which, per nucleon, neutron production is not high enough to bring on r-process. In this case they propose that neutron addition to Si^{28} follows the sequence $Si^{28}(n,\gamma)Si^{29}(n,\gamma)Si^{30}(n,\gamma)Si^{31}(\beta^-, 4\ \text{sec})P^{31}(n,\gamma)\cdot$ $\cdot P^{32}(\beta^-\ 21\ \text{days})S^{32}(n,\gamma)S^{33}(n,\alpha)Si^{30}$, explaining thereby the excess P concentration in χ Cnc, and predicting excess concentration of Si^{30} in stars with P excess. In χ Cnc, indeed, lines of SiII show strengthening; also in 3 Cen A, Bidelman has identified PII lines, and Si is overabundant (45). But, the reverse is not true; there are many stars with Si strong, but in which PII lines are not detectable. For example, the typical Mn-stars μ Lep and υ Her have considerably greater excess of Si, but λ4475, etc., lines of PII are not observed (these stars do not show rare-earth lines and therefore, presumably, r-process did not take place). Furthermore, if the above sequence does operate in χ Cnc and 3 Cen A, we should observe enhancement of SII lines also. But sulpher is underabundant in 3 Cen A to less than one-tenth (45). In χ Cnc, sulpher abundance has not been investigated.

(iv) For production of excess Mn, Fowler et al (36) have proposed that reactions

$$Fe^{54}(p,\gamma)Co^{55}(\beta^{+})Fe^{55}(\beta^{+})Mn^{55}$$

take place after the "flash". For the normal abundance we have, by number $\frac{Fe^{54}}{Fe} = .06$, $\frac{Mn^{55}}{Mn} = 1$, and $\frac{Fe^{54}}{Mn^{55}} = 5$ [or 1.7 according to Goldberg et al (in 22)]. Therefore, according to the above reactions, Mn can have, at the most, an overabundance by a factor of six, when all of the Fe^{54} is converted into Mn^{55} . This depletion of Fe^{54} should cause a spectroscopically imperceptible decrease (6 percent) to the total abundance of Fe. On the contrary, observationally, we find that in all the Mn-stars, Mn is "overabundant" by large factors (possibly up to 100 or more), and Fe is quite perceptibly under-abundant.

PECULIAR A-TYPE STARS AND EVOLUTION

Other authors (Eggen (33), Jascheks (43)) have described the various ideas about the evolutionary history of Ap stars. Here, only some spectroscopic characteristics will be point out which might have evolutionary significance: All the hotter Ap stars have Si normal or strong - none has Si weak. On the other hand, most of the cooler Ap stars have Si weak. Analogous results of Sargent and Searle (67), that oxygen is normal in the hotter Ap stars, but very weak in the cooler ones, may likewise have evolutionary significance.

Concentration of all the Mn-stars in a very narrow range of B-V color (Section II-B) is quite striking and suggests an effect of evolution. On the other hand, if Kraft's "yttrium star" HD22401 turns out to be an Mn-star, as we suspect (Section VI-A), it would rule out the hypothesis that Mn-stars are evolved, because HD22401 is a member of the young α Per cluster (Kraft (47)).

FURTHER WORK

The Ap-star studies so far may have provoked many new ideas and have provided valuable observational facts and constraints that must be satisfied by any theory or model. A great deal of extensive and concerted effort will be further required to reach a proper understanding. There are many lines of attack which suggest themselves, or which have been alluded to in the earlier discussion. As a result of the present study, these particularly interesting problems come to focus:

(a) Fewness of the broad-line Mn-stars is very conspicuous. We have discussed in Chapter II that this might be an effect of observational selection. Whether or not an Mn-group has only sharp-line stars has deep implications. Therefore, for the discovery of broad-line Mn-stars, a low dispersion survey of the ultraviolet (3400Å-3500Å) spectrum of B8-A0 stars should be made.

(b) Among "Sr"-stars, there exists a class in which Sr is strong but CaII-K is not weak. So far, in the literature, such stars have not been differentiated from those in which SrII lines are strong but CaII-K is very weak. In view of our findings that CaII-K is very weak in all the Cr-stars, it seems appropriate to make a distinction between the two types of Sr-stars.

A comparative study of the two types will provide valuable information.

To assist in the selection of Sr-stars in which the K-line is not weak for their temperature, a list of stars is given here. Since this list is based on the comments (some of which were ambiguous) encountered in the literature, a few of these may be spurious: (HD numbers) 10088, 15089, 15144, 16956, 42616, 45733, 78209, 118022, 148898, 190145, 207098, 213232, and 217401.

(c) The Ca-Ba correlation and the weakening of Si that we have found are likely to prove an important clue for an understanding of the phenomenon of Ap stars. Therefore, a more quantitative reinvestigation of these, with observational material of appropriate quality, will be very valuable.

(d) For a detailed abundance analysis, HD192913 is very strongly recommended. This is a fairly sharp-lined (ω = .2A) Si-4200 star with copious rare-earth and strong BeII lines.

The sharpline star γ Equ is a case of cooler star with Si strong. For this reason and because it can be used as an intermediate standard between the sun and the hotter Ap stars, an independent abundance redetermination for γ Equ is desirable.

TABLE VI-1

SUMMARY OF BEHAVIOR OF VARIOUS ELEMENTS IN AP STARS

	Mn	Si	Cr	Sr
Be	++, MS	++, ns	N, (+)	?
Mg	N, –	–	N, –	N, –
Si	N, +	++	+, N, –	–, N
Ca	N, (+)	N–*	––	–, (N)*
Sc*	+, (N)	+, –	+, –	+, –
Ti	±	N, (–)	N, –*	N, –*
V	(N, –)	?	(–, N)	(–)
Cr	N, –	±	++	+
Mn	++	N (–)	+, N	N (+)
Fe	–	+, (N)	N, (+)	(N)
Ni*	–	–	–	(–, N, +)
Ga*	+, (N)	?	?	?
Sr	+	+*	+	++
Y*	++		+, N, –	+; N
Zr*	+, N	+; N	+; N	+; N
Ba	nW*	?*	–*	–, nW
R. E. (Eu)	ns	+, ns	++, ns	++, (ns)

ns not 'strong'

nw not 'weak'

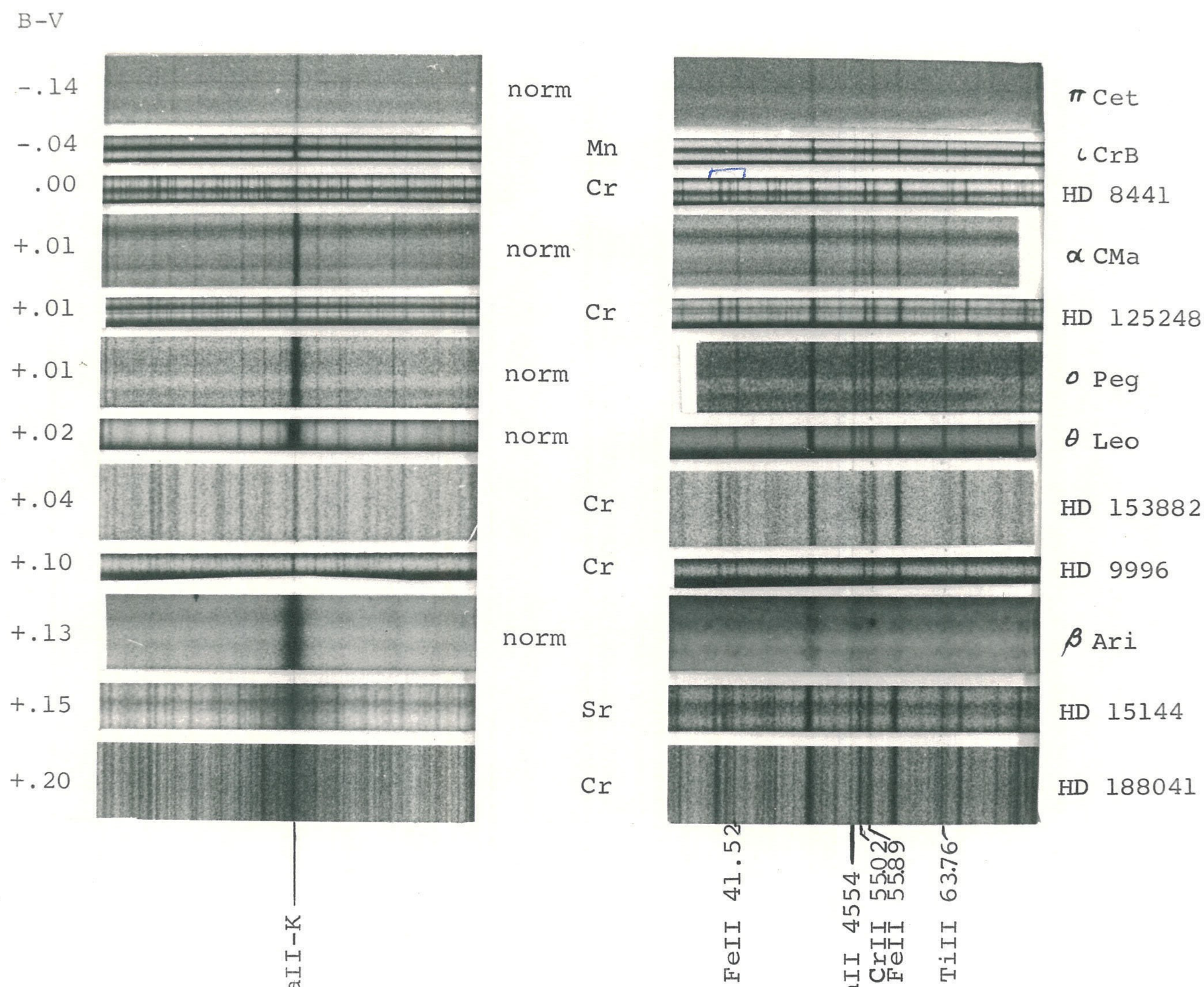
B-V
-.14
-.04
.00
+.01
+.01
+.01
+.02
+.04
+.10
+.13
+.15
+.20
norm
Mn
Cr
norm
Cr
norm
norm
Cr
Cr
norm
Sr
Cr
π Cet
ι CrB
HD 8441
α CMa
HD 125248
ο Peg
θ Leo
HD 153882
HD 9996
β Ari
HD 15144
HD 188041
CaII-K
FeII 41.52
BaII 4554
CrII 55.02
FeII 55.89
TiII 63.76

Figure VI-1

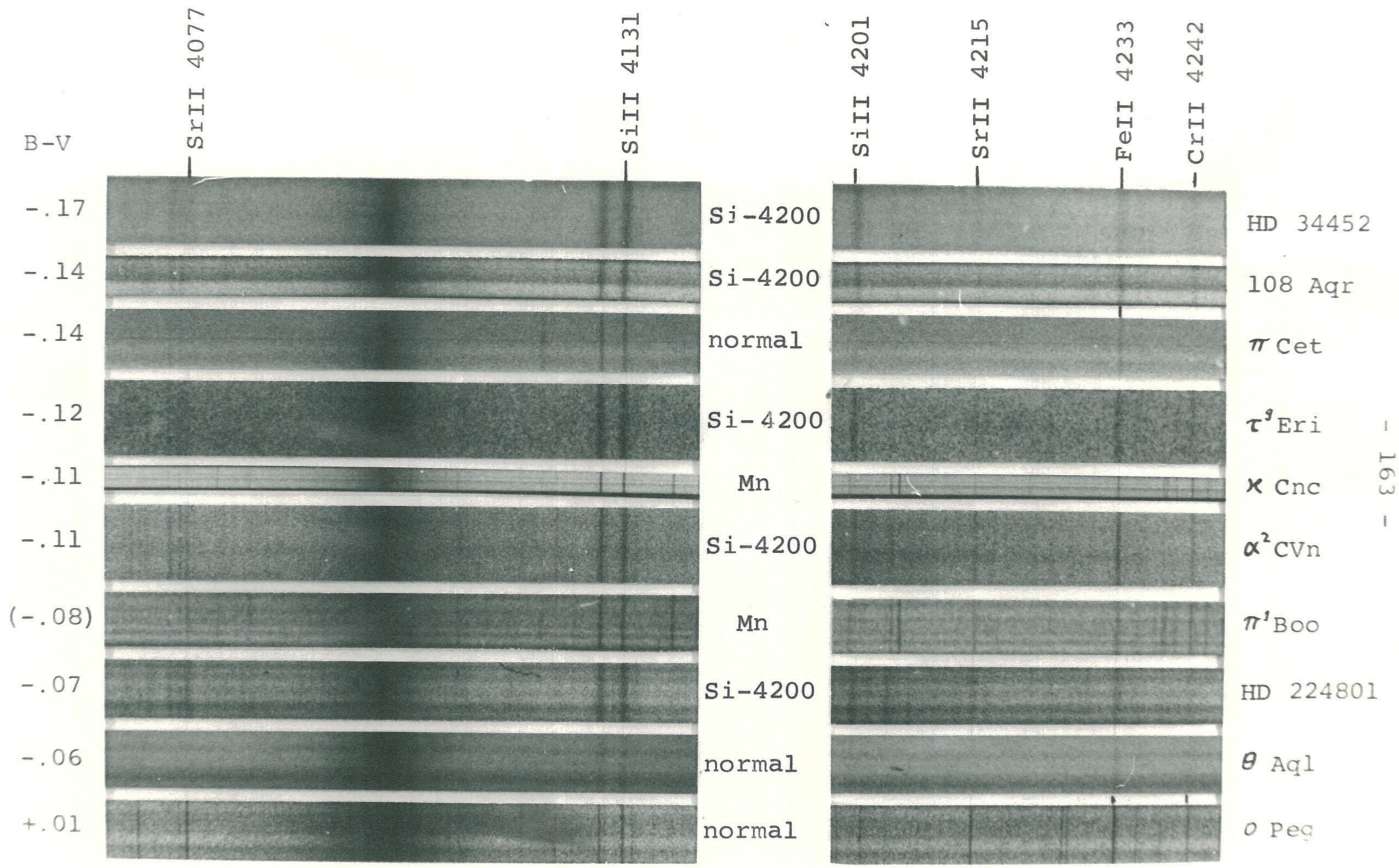
B-V
SrII 4077
SiII 4131
SiII 4201
SrII 4215
FeII 4233
CrII 4242
-.17
Si-4200
HD 34452
-.14
Si-4200
108 Aqr
-.14
normal
π Cet
-.12
Si-4200
τ³ Eri
-.11
Mn
κ Cnc
-.11
Si-4200
α² CVn
(-.08)
Mn
π¹ Boo
-.07
Si-4200
HD 224801
-.06
normal
θ Aql
+.01
normal
ο Peg

Figure VI-2

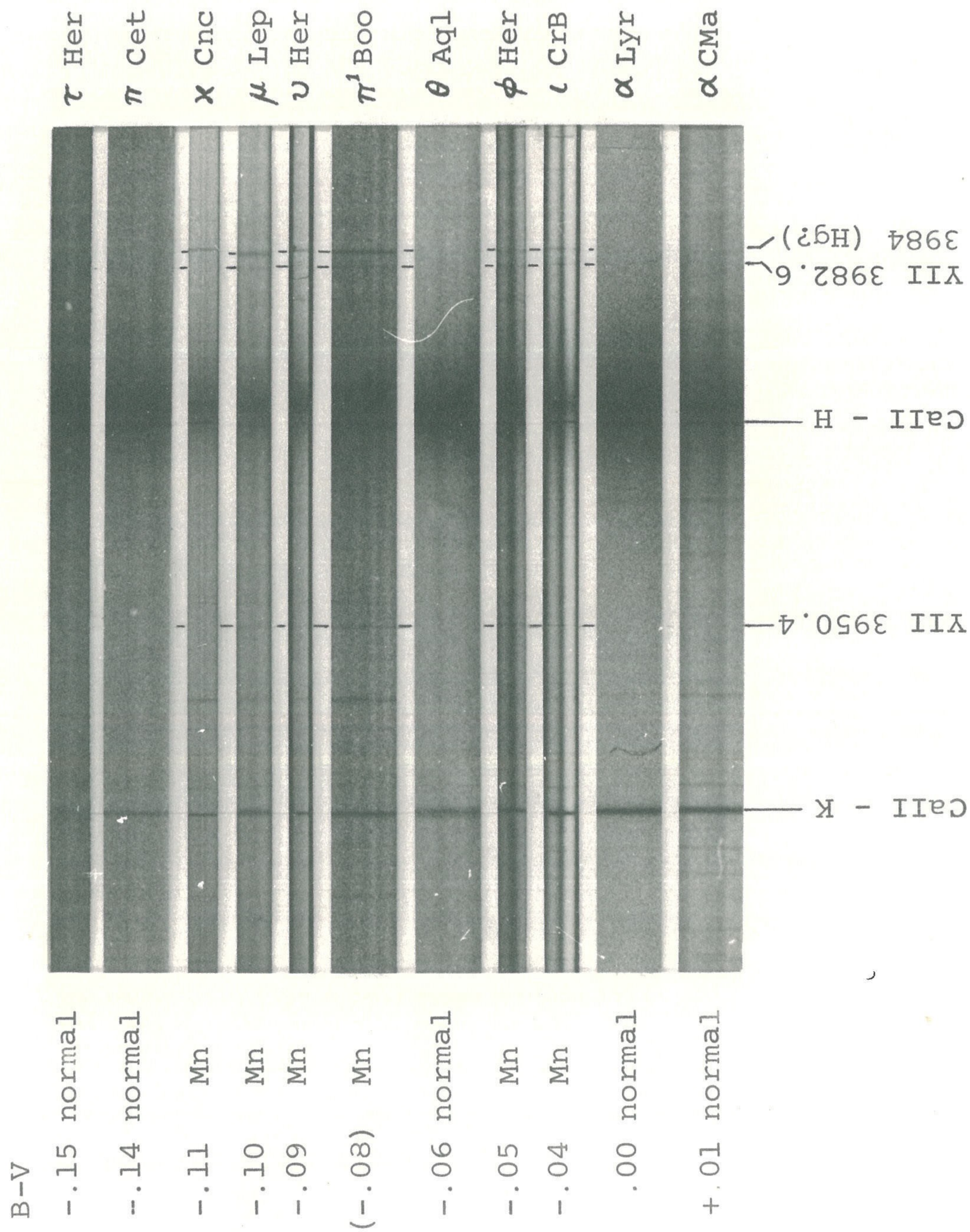
τ Her
π Cet
χ Cnc
μ Lep
υ Her
π¹ Boo
θ Aql
φ Her
ι CrB
α Lyr
α CMa
3984 (Hg?)
YII 3982.6
CaII - H
YII 3950.4
CaII - K
B-V
-.15 normal
-.14 normal
-.11 Mn
-.10 Mn
-.09 Mn
(-.08) Mn
-.06 normal
-.05 Mn
-.04 Mn
.00 normal
+.01 normal

Figure VI-3

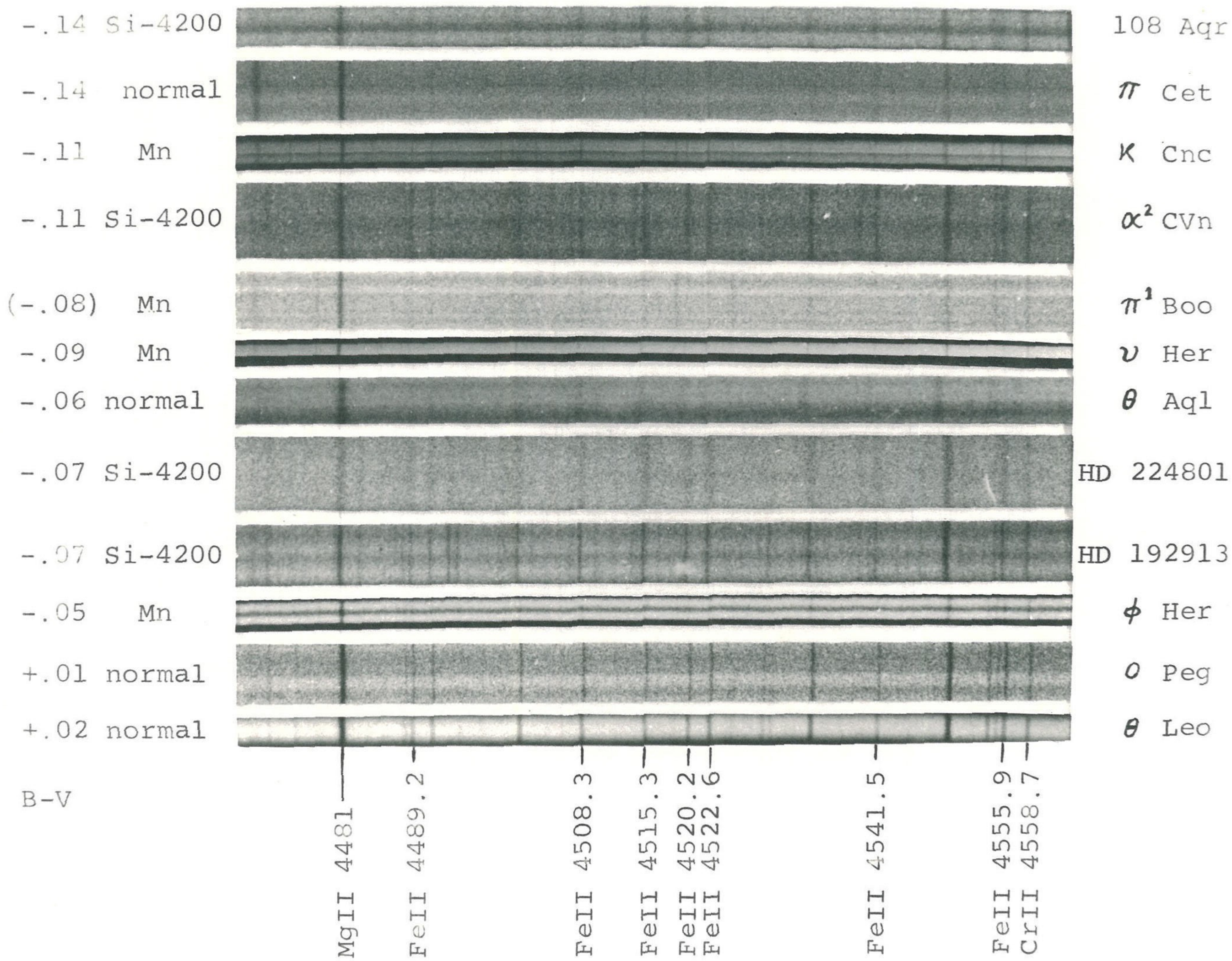
-.14 Si-4200
-.14 normal
-.11 Mn
-.11 Si-4200
(-.08) Mn
-.09 Mn
-.06 normal
-.07 Si-4200
-.07 Si-4200
-.05 Mn
+.01 normal
+.02 normal
B-V
108 Aqr
π Cet
κ Cnc
α^2 CVn
π^1 Boo
υ Her
θ Aql
HD 224801
HD 192913
ϕ Her
o Peg
θ Leo
MgII 4481
FeII 4489.2
FeII 4508.3
FeII 4515.3
FeII 4520.2
FeII 4522.6
FeII 4541.5
FeII 4555.9
CrII 4558.7

Figure VI-4

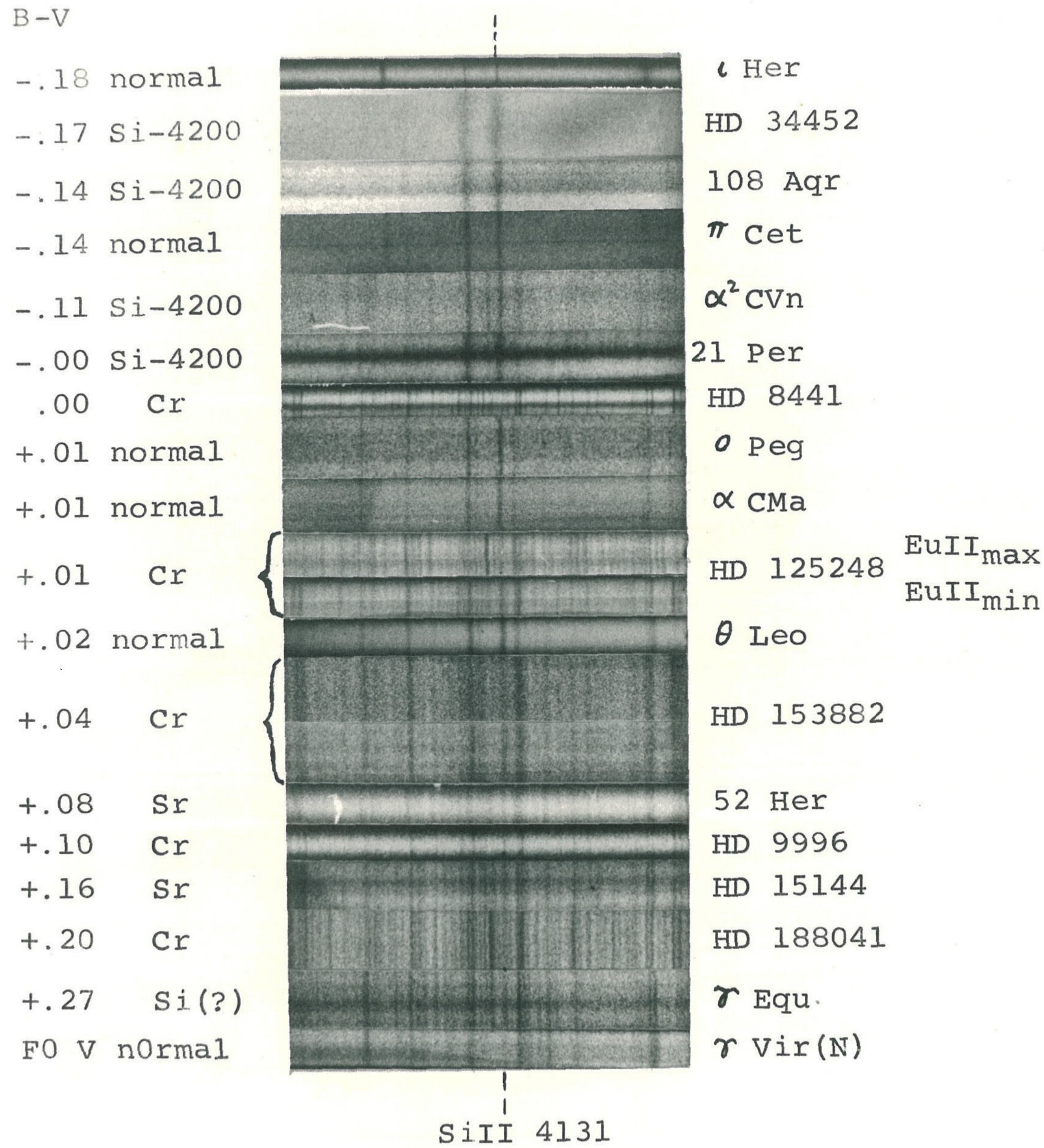

Figure VI-5

Stars with silicon 'strong', and 'weak'.

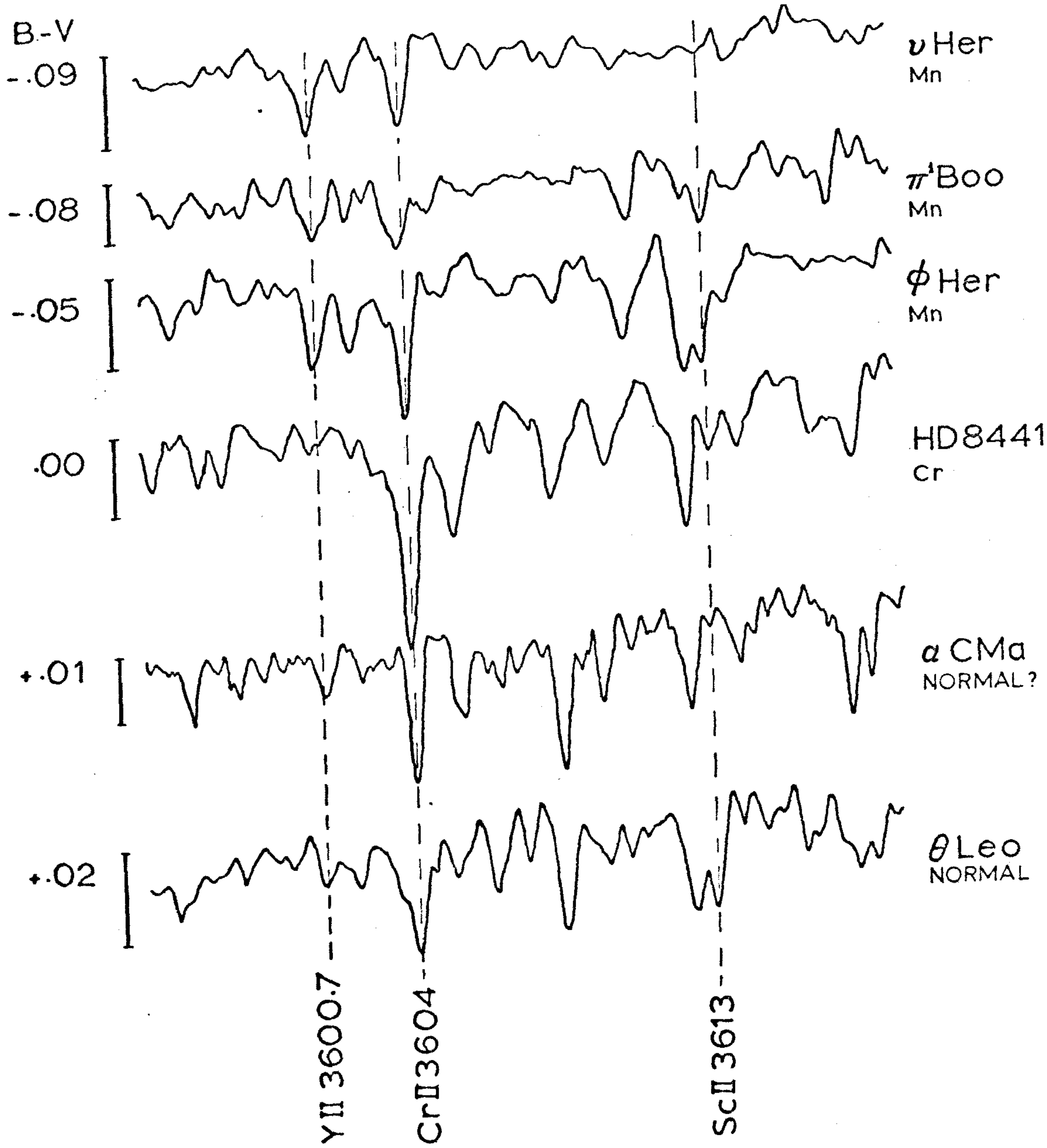

Figure VI-6

In Figs. VI-6,7&8 the vertical bars represent 10% of the continuum

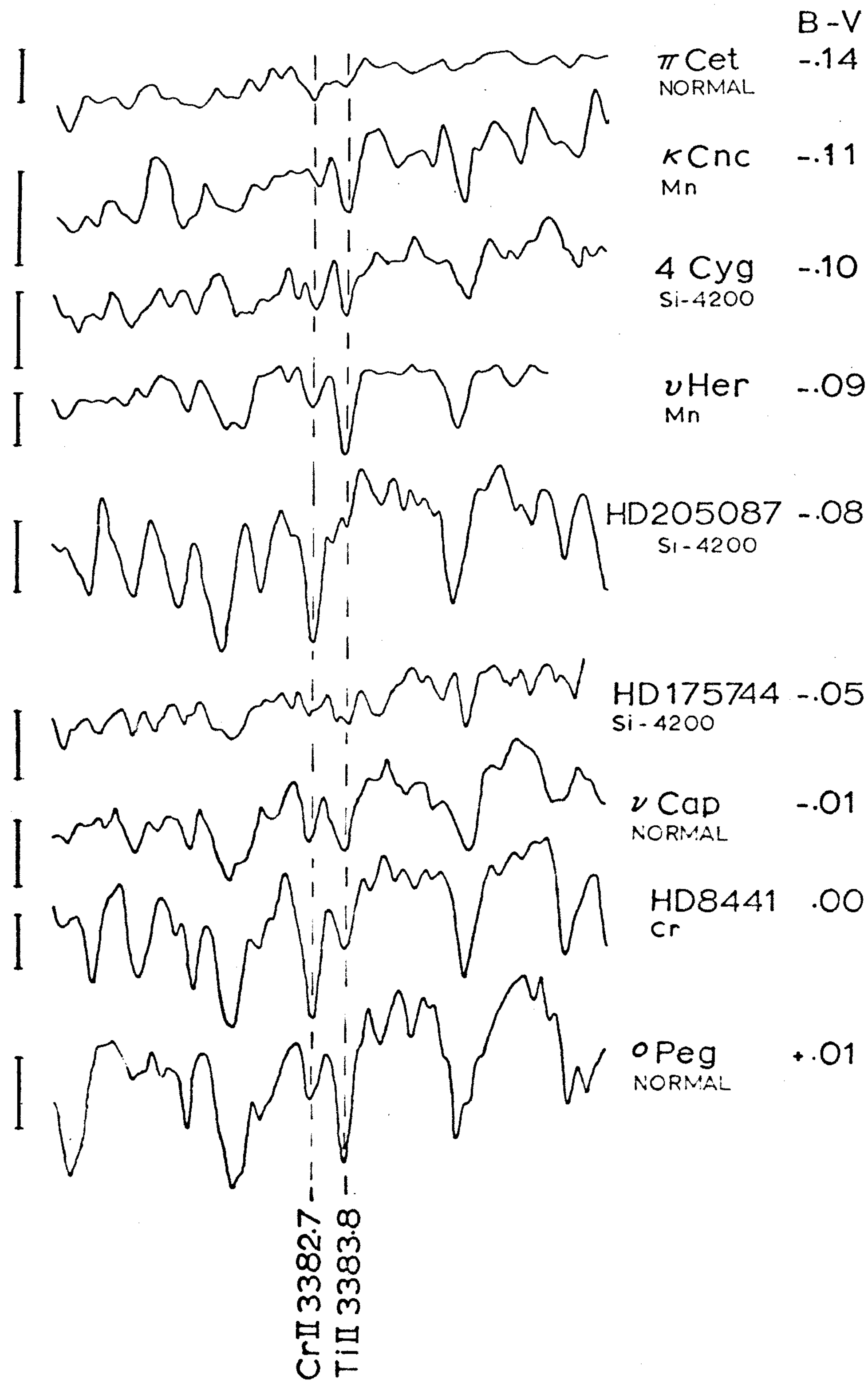
B-V
π Cet -.14
NORMAL
κ Cnc -.11
Mn
4 Cyg -.10
Si-4200
υ Her -.09
Mn
HD205087 -.08
Si-4200
HD175744 -.05
Si-4200
ν Cap -.01
NORMAL
HD8441 .00
Cr
o Peg +.01
NORMAL
CrII 3382.7
TiII 3383.8

Figure VI-7

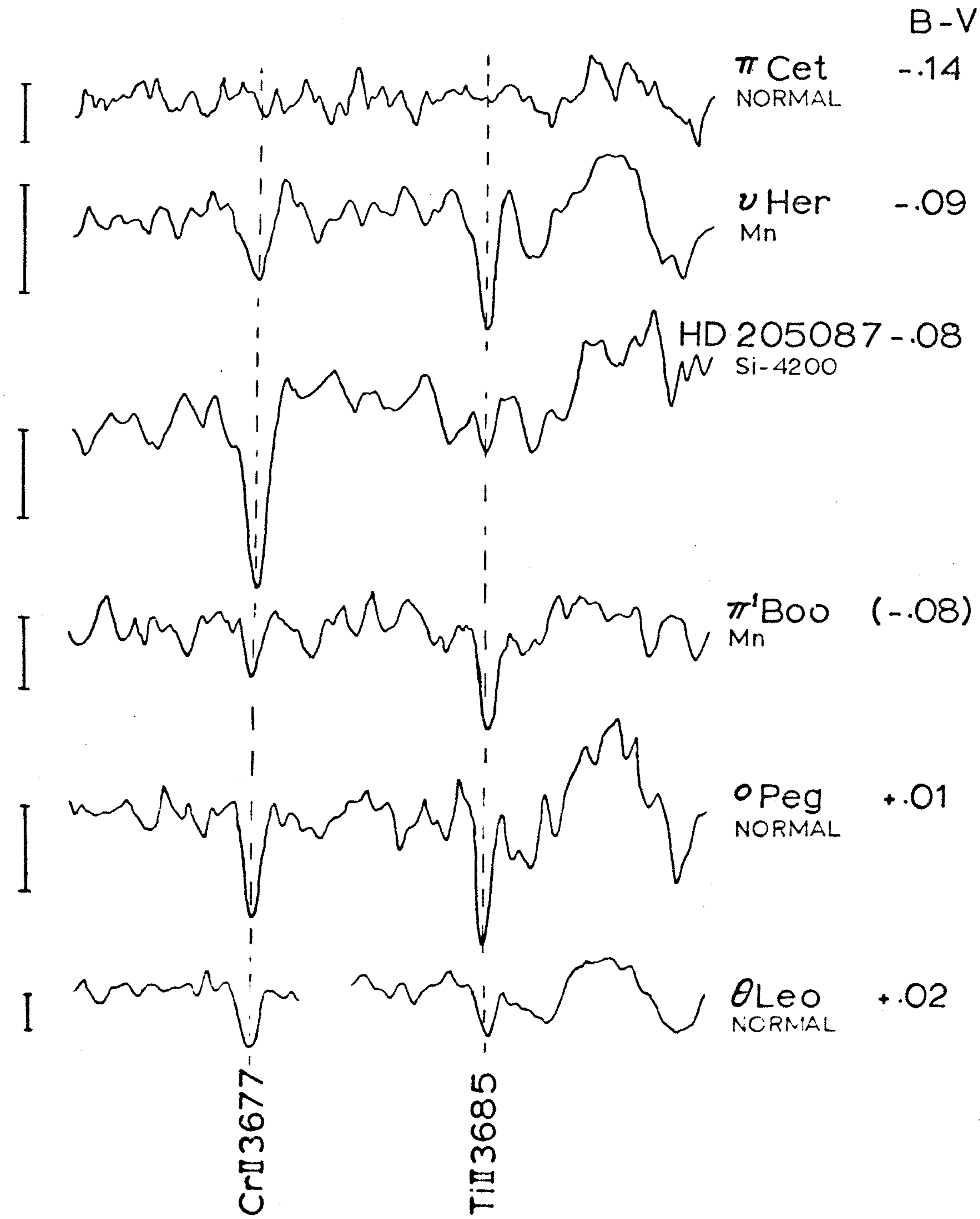
B-V
π Cet -.14
NORMAL
υ Her -.09
Mn
HD 205087 -.08
Si-4200
π¹ Boo (-.08)
Mn
o Peg +.01
NORMAL
θ Leo +.02
NORMAL
CrII 3677
TiII 3685

Figure VI-8

Symbols used in Figures VI-9 to VI-15:

■	(filled)	'Mn' stars	(Table III-4a)
□	(unfilled)	'Si' stars	(Table III-4b)
○		'Cr' stars	(Table III-4c)
▽		'Sr' stars	(Table III-4d)

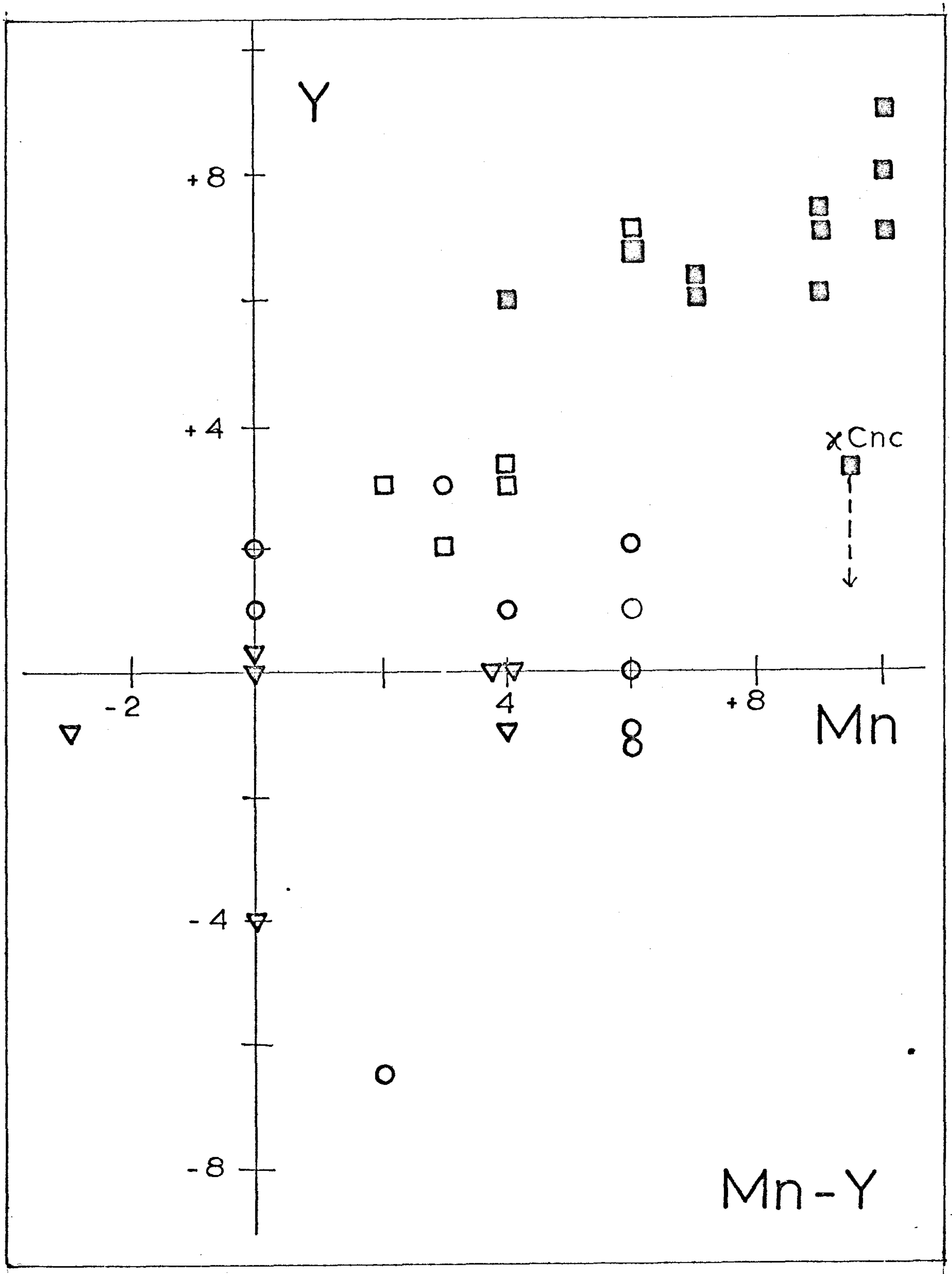

Figure VI-9

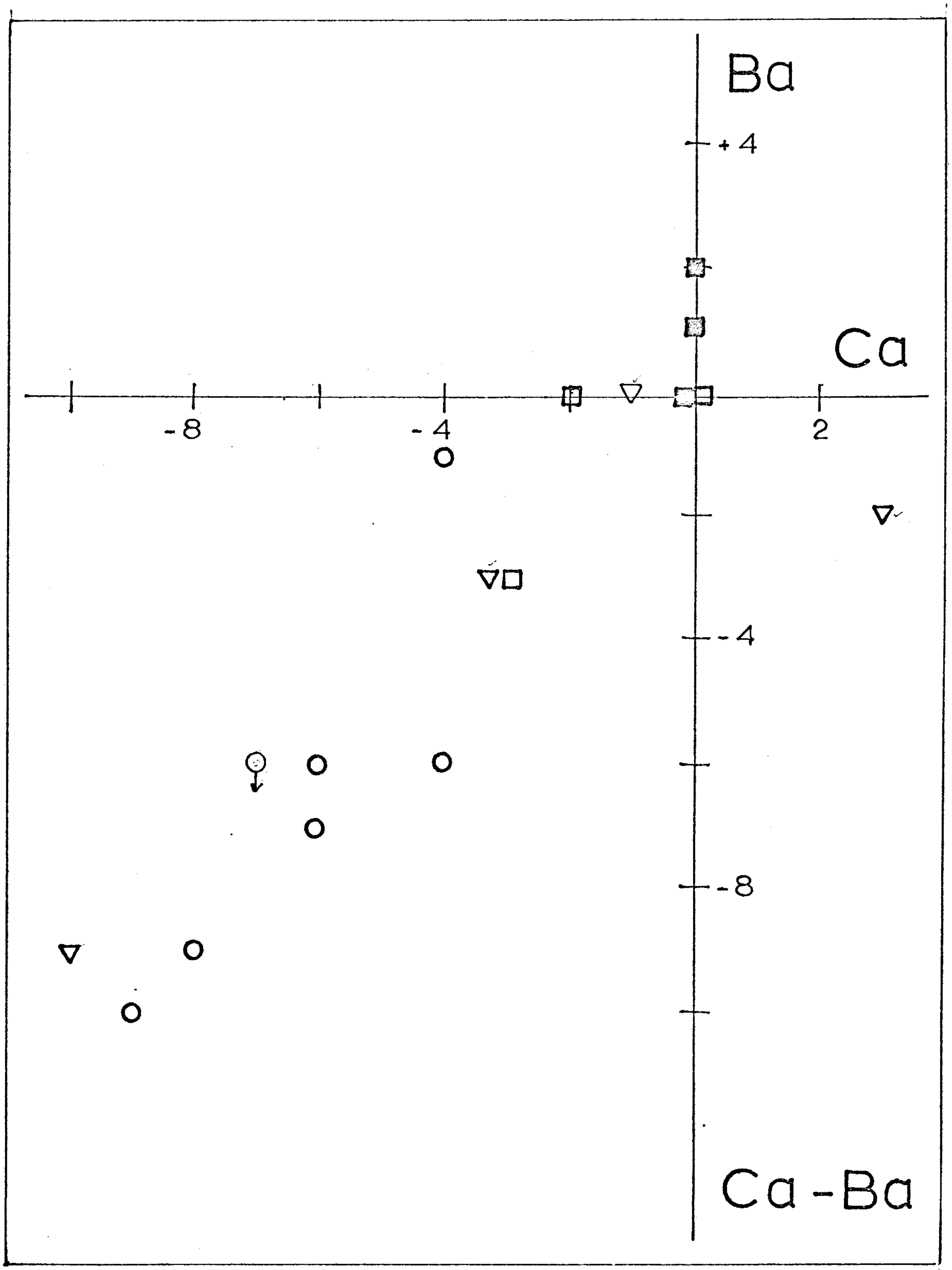

Figure VI-10

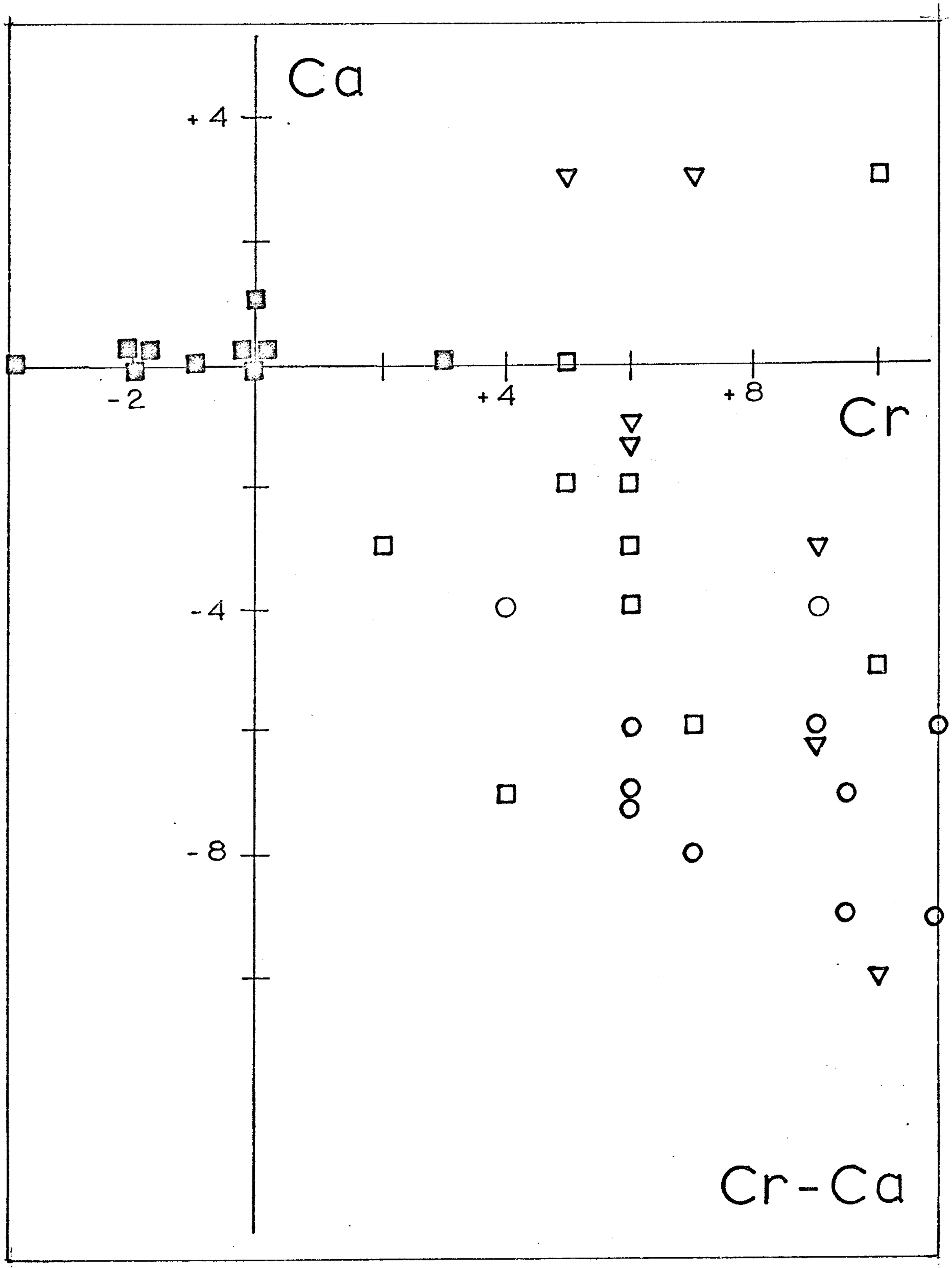

Figure VI-11

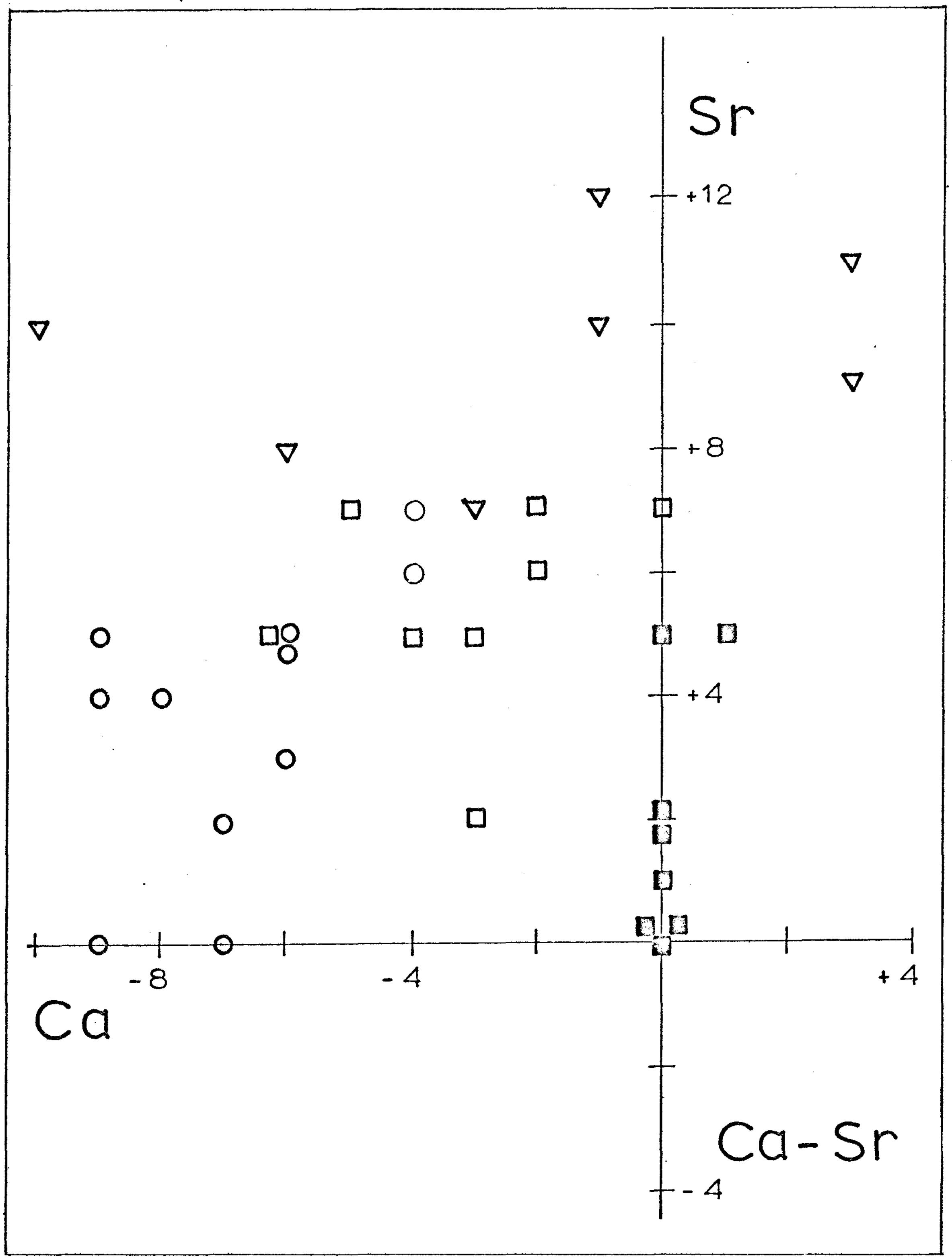

Figure VI-12

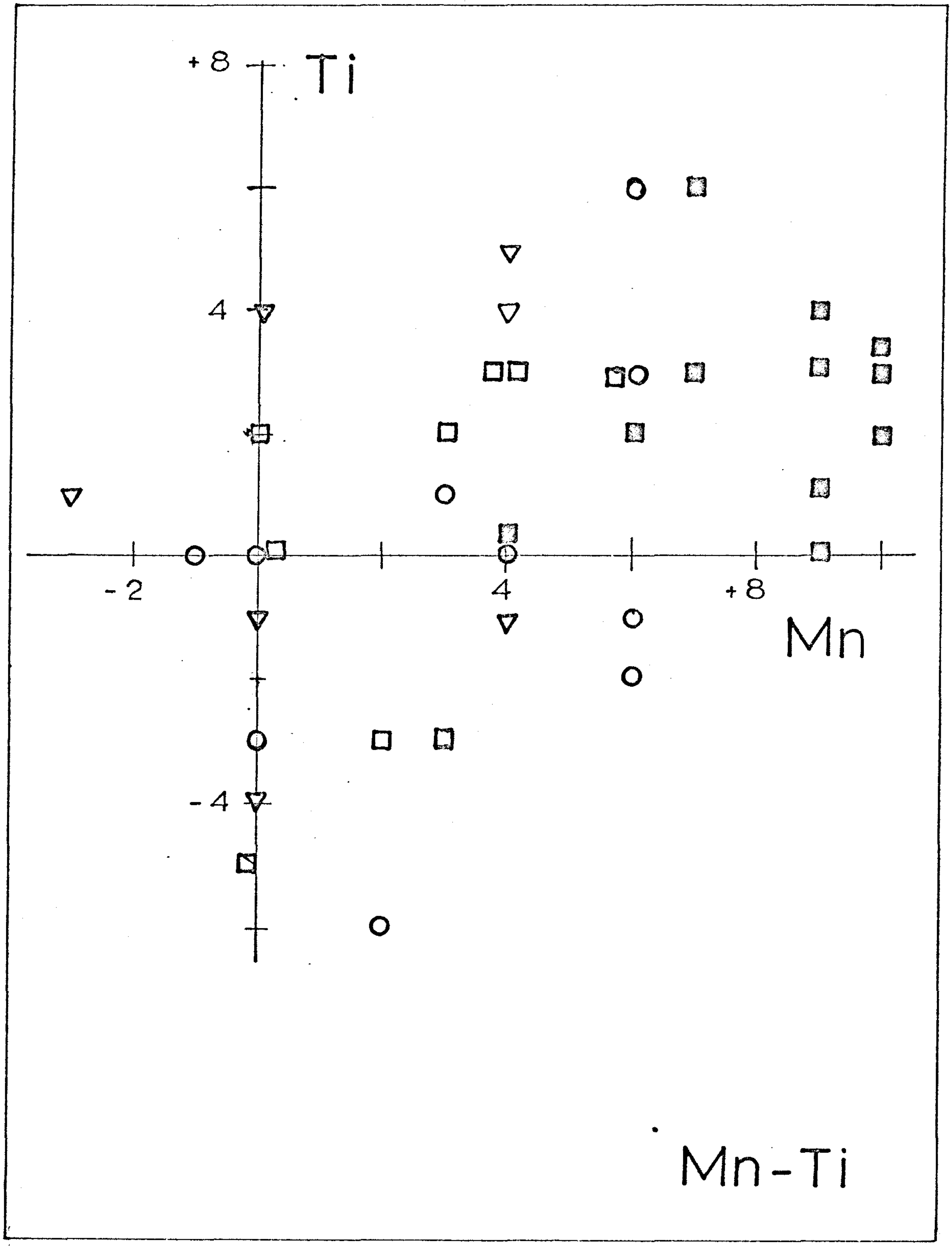

Figure VI-13

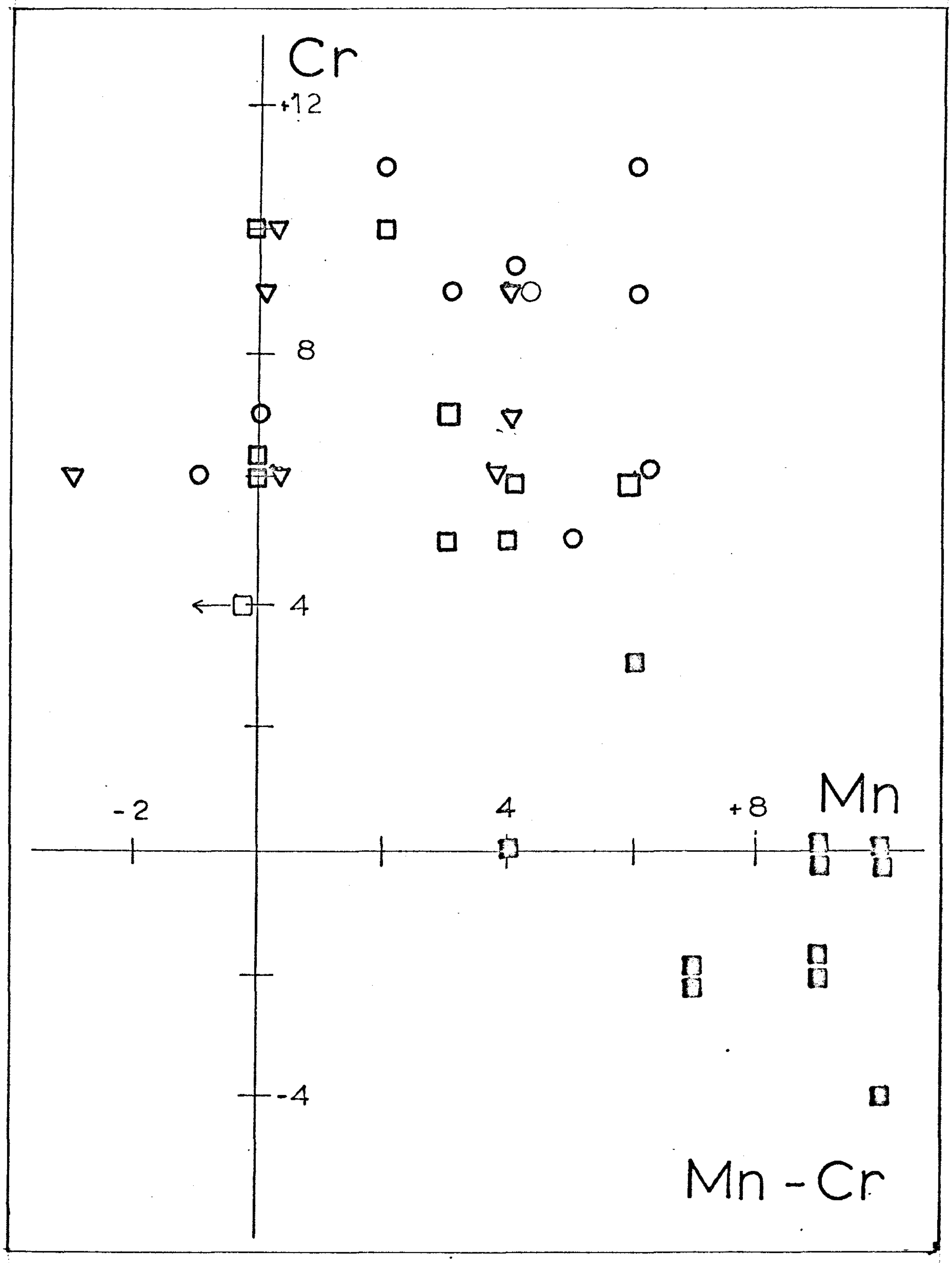

Figure VI-14

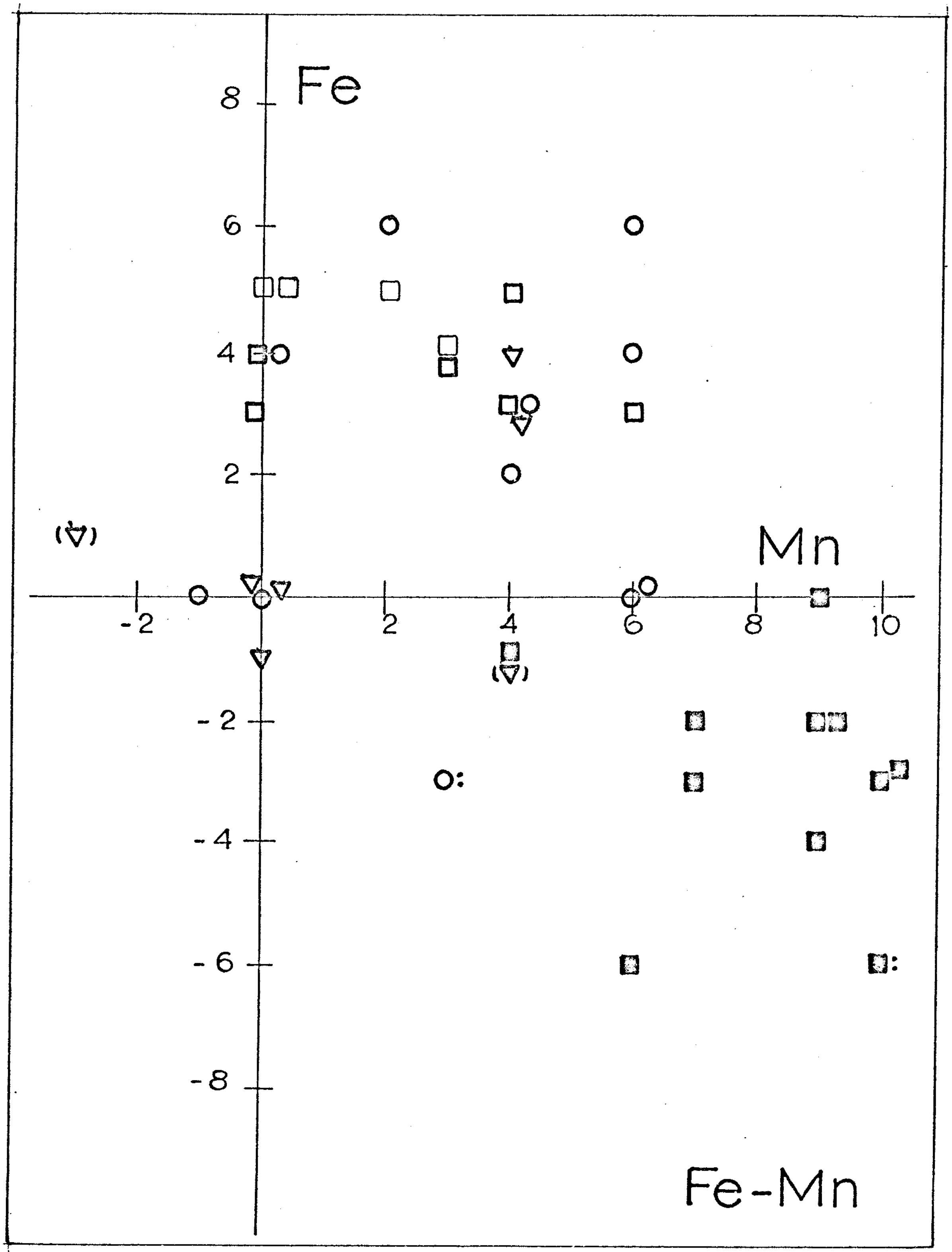

Figure VI-15

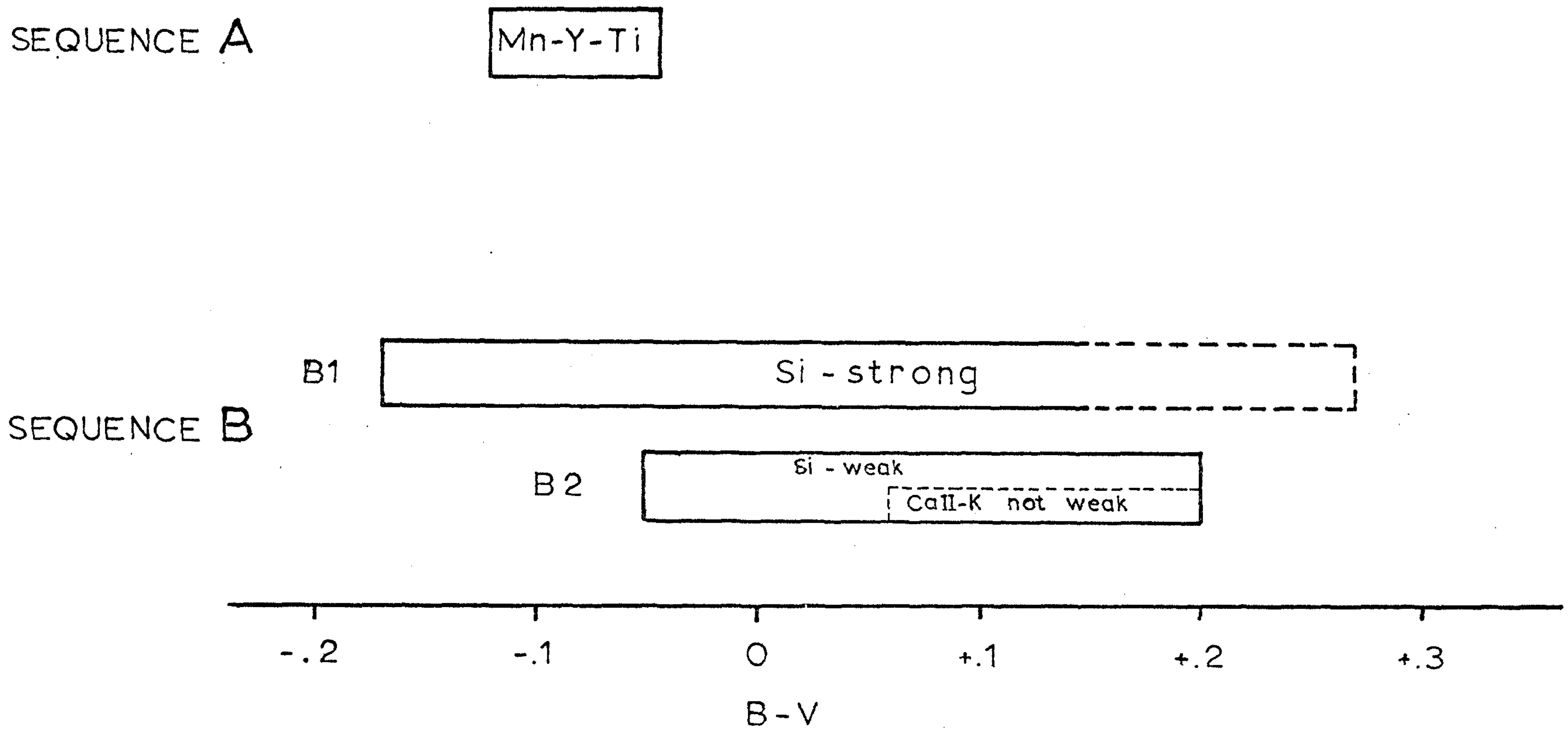
SEQUENCE A
Mn-Y-Ti
SEQUENCE B
B1
Si - strong
B2
Si - weak
CaII-K not weak
-.2
-.1
0
+.1
+.2
+.3
B-V

Figure VI-16

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APPENDIX I

Space Motion:

The components (Km/sec) of the space motion of a star are derived from the formulae :

$$X = 4.74 \frac{\mu_l}{p} \cos b \sin l + 4.74 \frac{\mu_b}{p} \cos l \sin b - V_r \cos l \cos b ,$$

$$Y = 4.74 \frac{\mu_l}{p} \cos b \cos l + 4.74 \frac{\mu_b}{p} \sin l \sin b - V_r \sin l \cos b ,$$

$$Z = 4.74 \frac{\mu_b}{p} \cos b + V_r \sin b ;$$

x-axis points towards $l^{II} = 0$, $b^{II} = 0$; Y-axis towards $l^{II} = 90$, $b = 0°$; and Z-axis towards $b^{II} = 90°$; and where is the galactic longitude, b is the galactic latitude, V_r is the radial velocity (Km/sec), p is the trigonometric parallax (in seconds of arc); and μ_l and μ_b are the l and b components of proper motion (seconds of arc per year).

In the Bright Star Catalogue (40), the right ascension and declination components (μ_α and μ_δ) of proper motion are given. These can be converted into μ_l and μ_b values by these formulae:

$$\mu_l \cos b = \mu_\alpha \cos \delta \cos \phi + \mu_\delta \sin \phi$$

$$\mu_b = -\mu_\alpha \cos \delta + \mu_\delta \cos \phi$$

where ϕ is given by

$$\sin\phi\cos b = \cos D\sin(\alpha - A)$$
$$\cos\phi\cos b = \sin D\cos\delta - \cos D\sin\delta\cos(\alpha - A)$$

where A and D , and α and δ are the right ascension and declination of the north galactic pole, and the star, respectively. F the new system (16) , $A = 12^h\ 49^m$ (1950) and $D = +27^\circ.4$ (1950).

APPENDIX II

LIST OF SPECTROGRAMS USED IN THIS STUDY

B-V	Star	Plate Number
-.18	<u>ι Her</u>	Ce17134 (UA); Ce3241 (b, 2.8)
-.17	34452	Ce15064 (b, B); Ce8239 (b, B); Pd 1223
-.15	<u>τ Her</u>	Ce17133 (UA); Ce9293 (b, B)
-.14	108Aqr	Ce15754 (UA); Ce15688 (UA); Ce11505 (bB); Ce11532 (b, B); Pb1091
-.14	<u>πCet</u>	Ce15749 (U, A); Ce15750 (U, A); Ce15765 (u+b, B)
-.12	τ^9Eri	Ce15684 (U, A); Pd1777, Pd1789, Pd1798
-.11	χ Cnc	(Ce16233, Ce16234) (U, A); Pb2531, Pb6574
-.11	124224	Ce15571 (b, B)
-.11	α^2CVn	Ce17126 (U, A); Ce17127 (U, A); Ce15570 (U+b; B); Ce15585 (U+b; B); Ce15594 (U+b; B)
-.10	α And	Xe7174 (overexposed side); Ce762(b) (prismatic spectrum average dispersion between $H\beta$ and H_{11} 10Mm)
-.10	172044	(Ce16230, Ce16231, Ce17135) (U, A); Pc5867 (b, overexposed)
-.10	μ Lep	Ce4416 (b, B); Ce5035 (b, 2.8)
-.10	4Cyg	[Ce15756, Ce15757, Ce16294] (U, A); Ce1555 Ce15558 (b; B)
-.09	112Her	Ce17129 (U+b, A)
-.09	υ Her	(Ce15677, Ce16287) (U, A), Ce14428 (b, 32^{11} + 133B)
	3322	Ce15746 (U, A); Xe7157
-.08	θAql	Ce15598 (U+b, B)
-.08	205067	[Ce15744, Ce15753] (U, A); Xe7149 (b)

B-V	Star	Plate Number
(-.08)	π^1 Boo	(Ce16226, Ce16286) (U,A); Ce15595 (b,B); Ce15596 (U+b;B); Pb2665
-.07	224801	(Ce15758, Ce15759) (U,A); (Ce15575, Ce15590) (U+b,B)
-.07	192913	Ce15745 (U,A); Ce15582 (U+b), Ce15592 (U+b), Ce17137 (a) U B
-.07	87Psc	Ce15747 (U,A)
-.05	175744	Ce16292 (U,A); Pc5868; Xe7154
-.05	ϕ Her	(Ce15685, Ce16288) (U,A); Pc3169
-.04	ι CrB	(Ce16244, Ce16245, Ce16283) (U,A); Pb 2556, Pb5695
-.04	μ Ser	Ce16290 (U,A)
-.04	17ComA	Ce17128 (U,A)
-.04	10783	(Ce15682, Ce15763) (U,A); Xe7158
(-.01)	η Leo	Ce17131 (U,A); Ce17132 (U,A)
(-.03)	ρ Her (A)	(Ce16228, Ce16229) (U,A)
-.01	ν Cap	Ce16290 (U,A)
-.00	21Per	(Ce15683, Ce15764) (U,A); Ce4885, Ce9040, Ce9094 (b,B)
.00	α Lyr	Ce15743 (U,A), Ce16232 (U,A); Xe7183 (b); Ce17137 (U,B)
.00	8441	(Ce15681, Ce15761) (U,A); Pb1559
+.01	α CMa	Ce15693 (U,A); Ce15766 (U+b,B)
+.01	o Peg	Ce15600 (U+b,B); Ce16237 (U,A); Ce17130 (U,A)
+.01	125248	(Ce5149, Ce5627) (b,4.5); Ce15593 (U,B); Pb5816
$+.01_5$	4778	Ce15748 (U,A); Xe7168

B-V	Star	Plate Number
+.02	θ Leo	Ce16242 (U,A), Ce16289 (U,A), Ce4249 (b,2.8); Ce4153 (b,B)
+.04	107612	(Ce16240, Ce16241) (U,A)
+.04	153882	(Ce15572, Ce15587, Ce15588, Ce15597) (U+b,B)
	γ Ari (s)	Ce15575 (U+b,B)
+.05	21Com	Xe6992; Pe1996; (Ce16238, Ce16239) (U,A)
$+.06_5$	2453	(Ce15680, Ce15755) (U,A); Xe7167, Xe7187
+.08	52Her	(Ce15751, Ce15752, Ce16227) (U,A); Ce8072 (b,4.5)
+.09	204411	(Ce15678, Ce16293) (U,A); Pb6156
+.10	9996	(Ce15689, Ce15762) (U,A); Ce5413 (b,B); Xe7156
+.13	9Tau	Ce15692 (U,B); Xe7151
+.14	β Ari	Ce15050 (U+b,B), Ce15690 (U,A), Ce15852 (b,4.5)
+.16	15144	Ce15691 (U,A); (Ce9039, Ce9049) (b,B)
+.16	72Oph	Ce16284 (U,A); Ce16285 (U,A)
+.16	21LMi	Ce16225 (U,A)
+.16	164258	(Ce16235, Ce16236, 16291) (U,A)
+.17	ι UMa	Ce16223 (U,A), Ce16224 (U,A)
+.19	λPsc	Ce16352 (b,B)
+.20	188041	(Ce15573, Ce15589) (b,B), Xe7163
+.22	191742	(Ce15686, Ce15760) (U,A)
+.27	γ Equ	Ce15679 (U,A); Ce15585 (b,B); Ce15601 (U+b,B)
(FOV)	γ Vir(N)	Ce7990 (b,B); Ce8052 (b,B)

Remarks

1. Ap and normal stars are arranged in order of their B-V value. Normal stars are indicated by an underscore.

2. Notation:

U	Ultraviolet plate	
b	Blue plate	
A	21A/mm dispersion	
B	10 A/mm dispersion	
Xe	40 A/mm dispersion	60-inch plates
Xd	20 A/mm dispersion	
Pb	4.5 A/mm dispersion	Palomar plates only for blue
Pc	9 A/mm dispersion	
Pd	18 A/mm dispersion	
Ce	100-inch plates	
2.8, 4.5, etc.	dispersion of the plate in A/mm	

3. Star numbers refer to HD Catalog.