Heavy element nucleosynthesis: the r process

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HELMHOLTZ RESEARCH FOR GRAND CHALLENGES

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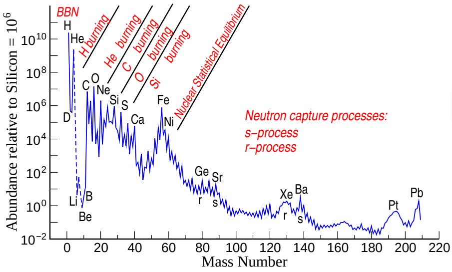
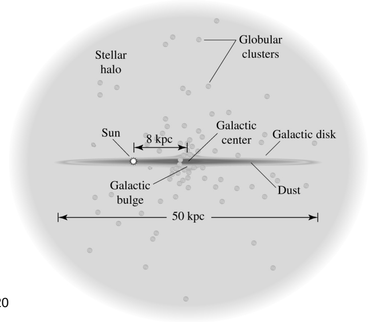
Outline

Solar system abundances

- Solar system abudances contain signatures of nuclear structure and nuclear stability.
- They are the result of different nucleosynthesis processes operating in different astrophysical environments and the chemical evolution of the galaxy.

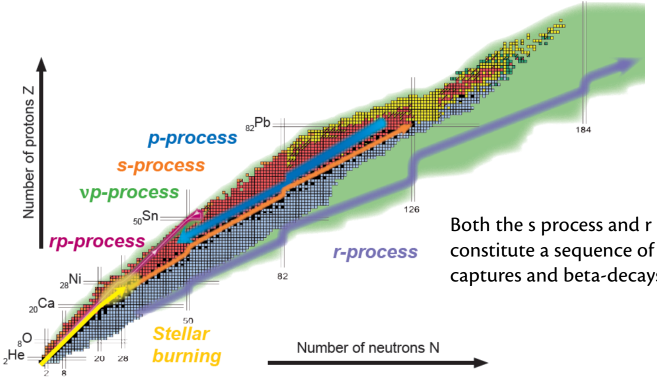




K. Lodders, Astrophys. J. **591**, 1220-1247 (2003)

Nucleosynthesis processes

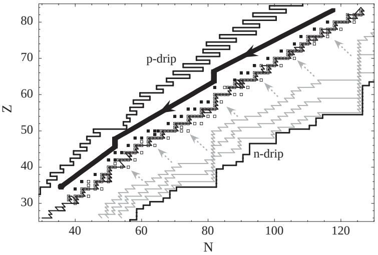
In 1957 Burbidge, Burbidge, Fowler and Hoyle and independently Cameron, suggested several nucleosynthesis processes to explain the origin of the elements.

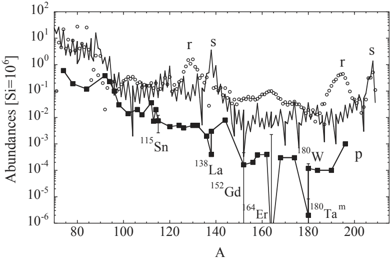


Both the s process and r process constitute a sequence of neutron captures and beta-decays.

Nucleosynthesis beyond iron

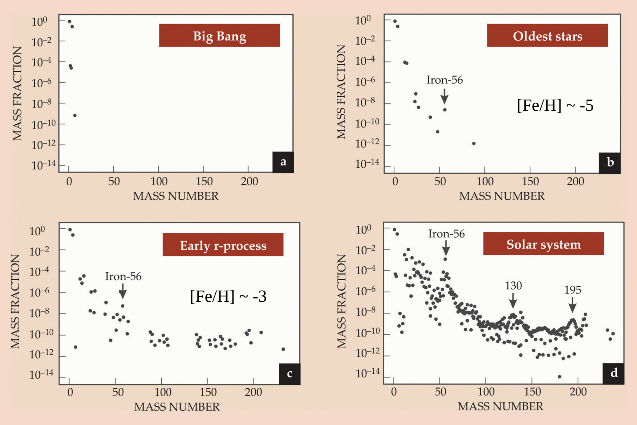
Three processes contribute to the nucleosynthesis beyond iron: s process, r process and p process (γ-process).





- s process: low neutron densities, $n_n = 10^{10-12}$ cm^{-3}, $\tau_n > \tau_\beta$ (site: intermediate mass stars)
- r process: large neutron densities, $n_n > 10^{20}$ cm^{-3}, $\tau_n \ll \tau_\beta$ (unknown astrophysical site)
- The r-process abundance is obtained by subtracting the calculated s-process abundance from the observed solar abundance of heavy elements, $N_r = N_\odot - N_s$. Very different odd-even staggering in s and r abundances.
- Can we validate the above assumptions?

Time evolution: metalicity



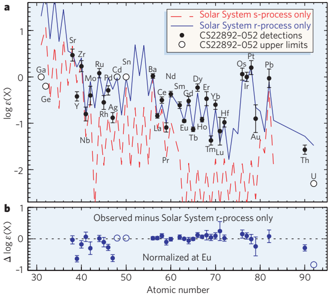
Astronomers use the metalicity:

$$[\mathrm{Fe/H}] = \log_{10}\left(\frac{N_{\mathrm{Fe}}}{N_{\mathrm{H}}}\right)_{*} - \log_{10}\left(\frac{N_{\mathrm{Fe}}}{N_{\mathrm{H}}}\right)_{\odot}$$

as a proxy for age.

Metal-poor star observations

- The r-process is known to occur already at low metalicities, early times, while the s-process occurs later in galactic history.
- If we observe old enough stars we may be able to see a star that contains only r-process and no s-process.
- This may also allow to observe the contribution of a single r-process event to the solar composition.

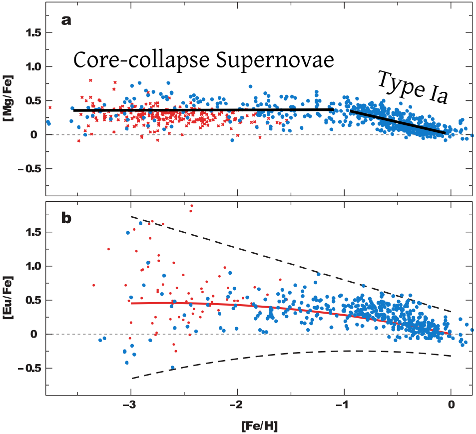


$$\log \varepsilon(X) = \log_{10}(N_X/N_H) + 12.0$$

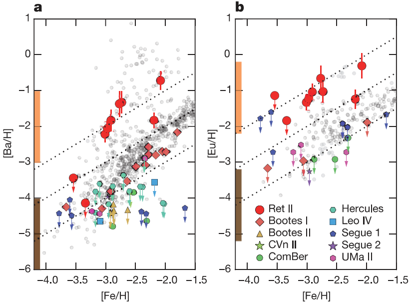
from Cowan and Sneden, Nature **440**, 1151 (2006)

Implications from observations

Individual stars, Milky Way Halo
Sneden, Cowan & Gallino, 2008



Ji et al 2016 found that only 1 of 10 ultrafaint dwarf galaxies is enriched in r-process elements

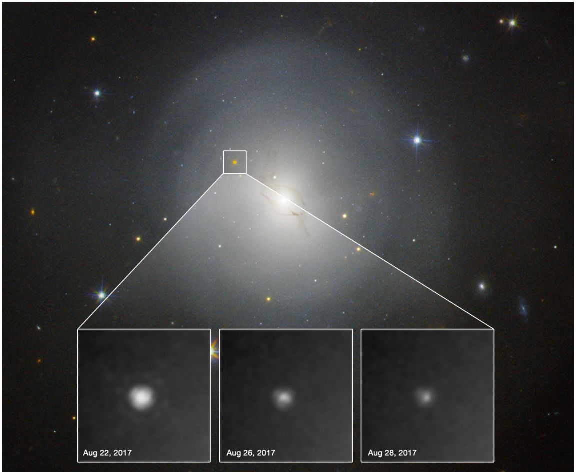


R process related to rare high yield events not correlated with Iron enrichment

Similar results obtained by ^{60}Fe and ^{244}Pu observations in deep sea sediments (Wallner et al, 2015; Hotokezaka et al, 2015)

Kilonova/Macronova luminosity

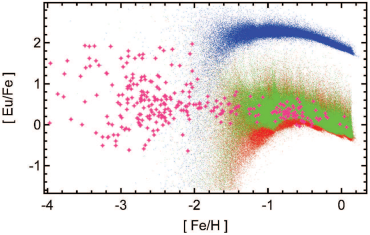
First direct signal from “in situ” r process operation.



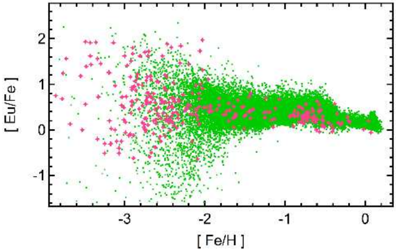
Credit: NASA & ESA. N. Tanvir (U. Leicester), A. Levan (U. Warwick), and A. Fruchter and O. Fox (STScI)

Introduction: Summary

- The r process is a primary process operating in a site that produces both neutrons and seeds. Large neutron densities imply a site with extreme conditions of temperature and/or density.
- There is strong evidence that the bulk of r-process content in the Galaxy originates from a high yield/low frequency events.
- Neutron star mergers may account for most of the r-process material in the galaxy. However, due to the coalescent delay time they may not contribute efficiently at low metalicities. Magneto-rotational supernova may contribute at low metalicities.



- Red dots: 10^8 yr coalescence time
- Green dots: 10^6 yr coalescence time
- Blue dots: larger merger probability.



- Including MHD-jet supernovae

Wehmeyer, B., M. Pignatari, and F.-K. Thielemann , Mon. Not. Roy. Astron. Soc. 452, 1970 (2015)

Types of reactions

Nuclei in the astrophysical environment can suffer different reactions:

- Decay (decay rate: $\lambda = \ln 2/t_{1/2}$)

$$^{56}\mathrm{Ni} \rightarrow {}^{56}\mathrm{Co} + e^{+} + \nu_e$$
$$^{15}\mathrm{O} + \gamma \rightarrow {}^{14}\mathrm{N} + p$$

$$\frac{dn_a}{dt} = -\lambda_a n_a$$

In order to dissentangle changes in the density (hydrodynamics) from changes in the composition (nuclear dynamics), the abundance is introduced:

$$Y_a = \frac{n_a}{n}, \quad n \approx \frac{\rho}{m_u} = \text{Number density of nucleons (constant)}$$

$$\frac{dY_a}{dt} = -\lambda_a Y_a$$

Rate can depend on temperature and density

Types of reactions

Nuclei in the astrophysical environment can suffer different reactions:

- Capture processes

$$a + b \rightarrow c + \gamma$$

$$\frac{dn_a}{dt} = -n_a n_b \langle \sigma v \rangle$$

$$\frac{dY_a}{dt} = -\frac{\rho}{m_u} Y_a Y_b \langle \sigma v \rangle$$

destruction rate particle a by reaction with b: $\lambda_a(b) = \rho Y_b \langle \sigma v \rangle / m_u$

- photodissociation rates

$$\gamma + c \rightarrow a + b$$

$$\frac{dY_c}{dt} = -Y_c \lambda_c = -Y_c n_\gamma \langle \sigma c \rangle$$

$\langle \sigma c \rangle$ photodissociation cross section averaged over thermal photon spectrum.

The balance between capture and photodissociation is determined by the photon-to-baryon ratio.

Three body reactions

Due to the fact that there is no stable nuclei with $A = 5$ and 8, nuclei heavier than ^{4}He have to be build by 3-body reactions. The most relevant reactions are:

3-α Dominant in proton-rich environments

$$3\,^4\mathrm{He} \rightarrow {}^{12}\mathrm{C} + \gamma$$

$$\frac{dY_\alpha}{dt} = -\frac{3}{3!} Y_\alpha^3 \left(\frac{\rho}{m_u}\right)^2 \langle\alpha\alpha\alpha\rangle$$

$\alpha\alpha n$ Dominant in neutron-rich environments

$$2\,^4\mathrm{He} + n \rightarrow {}^{9}\mathrm{Be} + \gamma$$

$$\frac{dY_\alpha}{dt} = -\frac{2}{2} Y_\alpha^2 Y_n \left(\frac{\rho}{m_u}\right)^2 \langle\alpha\alpha n\rangle$$

These reactions are key for the build-up of heavy nuclei.

Reaction rates

Consider n_a and n_b particles per volume of species a and b. The number of nuclear reactions per unit of time and volume

$$a + A \rightarrow B + b$$

is given by:

$$r_{aA} = \frac{n_a(v_a) n_A(v_A)}{(1 + \delta_{aA})} \sigma(v) v, \quad v = |\boldsymbol{v}_a - \boldsymbol{v}_A| \text{ (relative velocity)}$$

In stellar environment the velocity (energy) of particles follows a thermal distribution that depends of the type of particles.

- Nuclei (Maxwell-Boltzmann)

$$n(v)dv = n 4\pi v^2 \left(\frac{m}{2\pi kT}\right)^{3/2} \exp\left(-\frac{mv^2}{2kT}\right) dv = n\phi(v)dv$$

- Electrons, Neutrinos (if thermal) (Fermi-Dirac)

$$n(p)dp = \frac{g}{(2\pi\hbar)^3} \frac{4\pi p^2}{e^{(E(p)-\mu)/kT} + 1} dp$$

- photons (Bose-Einstein)

$$n(p)dp = \frac{2}{(2\pi\hbar)^3} \frac{4\pi p^2}{e^{pc/kT} - 1} dp$$

Stellar reaction rate

The product σv has to be averaged over the velocity distribution $\phi(v)$ (Maxwell-Boltzmann)

$$\langle \sigma v \rangle = \int_0^\infty \phi(v)\sigma(v) v dv$$

that gives:

$$\langle \sigma v \rangle = 4\pi \left(\frac{m}{2\pi kT}\right)^{3/2} \int_0^\infty v^3 \sigma(v) \exp\left(-\frac{mv^2}{2kT}\right) dv, \quad m = \frac{m_a m_b}{m_a + m_b}$$

or using $E = mv^2/2$

$$\langle \sigma v \rangle = \left(\frac{8}{\pi m}\right)^{1/2} \frac{1}{(kT)^{3/2}} \int_0^\infty \sigma(E) E \exp\left(-\frac{E}{k_B T}\right) dE \qquad (1)$$

Cross section determination

The calculation of the cross section requires the determination of the wave function for the system projectile (a) and target (A) for a particular value of energy E. This requires solutions of the Schrodinger equation for a potential

$$V(r) = V_{\text{nuclear}}(r) + V_{\text{coulomb}}(r) + V_{\text{centrifugal}}(r)$$

- Nuclear potential: complicated form with strong dependence on energy, E, angular momentum, J and parity, π (due to the internal structure of the target and projectile). It is of very short range: $R = 1.2(A_a^{1/3} + A_A^{1/3})$ fm.
- Coulomb potential (only for charged particles):

$$V(r) = \frac{Z_a Z_A e^2}{r}$$

- Centrifugal barrier:

$$V(r) = \frac{\hbar^2 l(l+1)}{2mr^2}$$

  cross section suppressed for high l values. Normally s-wave ($l = 0$) and p-wave ($l = 1$) dominate.

Cross section is mainly determined by long range behaviour of the potential

Cross section

The general form of the total cross section for the formation of a nucleus with $A_C = A_a + A_A$ and $Z_C = Z_a + Z_A$

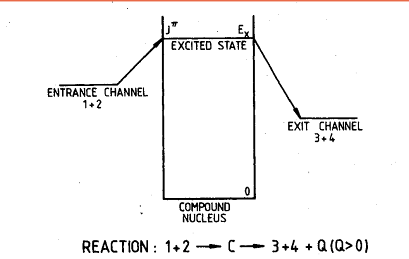
$$a + A \rightarrow C \rightarrow B + b$$

$$\sigma(E) = \pi\lambdabar^2 \sum_l (2l+1)T_l, \quad \lambdabar = \frac{\hbar}{mv} = \frac{\hbar}{\sqrt{2mE}}$$

T_l transmission coefficient through the potential barrier.

The problem reduces to a calculation of the tunneling probability through a barrier.

Neutron capture



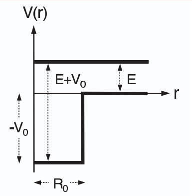


Fig. 2.7 Three-dimensional square-well potential of radius R_0 and potential depth V_0. The horizontal line indicates the total particle energy E. For the calculation of the transmission coefficient, it is necessary to consider a one-dimensional potential step that extends from $-\infty$ to $+\infty$. See the text.

$$A + n \rightarrow B + \gamma$$

$$\sigma_n(E) = \pi \lambda\!\!\!^{-2} \sum (2l+1) T_{l,n}(E) P_\gamma(E+Q)$$

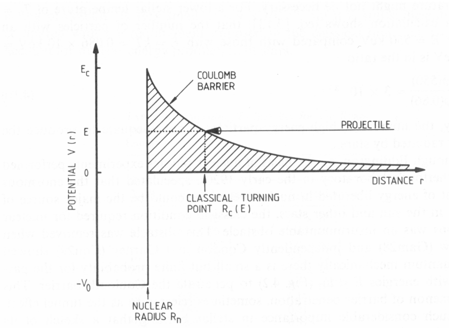
$T_{l,n}$ transmision coefficient, P_γ probability of gamma emission, E neutron energy ($\sim$ keV), $Q = m_A + m_n - m_B = S_n$, $Q \gg E_n$. Normally s-wave dominates and we have

$$\sigma_n(E) \propto \lambda\!\!\!^{-2} T_{0,n}(E), \quad T_n(E) \propto v$$

$$\sigma_n(E) \propto \frac{1}{v^2} v = \frac{1}{v}, \quad \langle \sigma_n v \rangle = \text{constant}$$

Charged-particle reactions

Stars' interior is a neutral plasma made of charged particles (nuclei and electrons). Nuclear reactions proceed by tunnel effect. For the $p + p$ reaction the Coulomb barrier is 550 keV, but the typical proton energy in the Sun is only 1.35 keV.



Assuming s-wave dominates:

$$\sigma(E) = \pi \lambda\!\!\!^{-2} T_0(E), \quad T_0 = \exp\left\{-\frac{2}{\hbar}\int_{R_n}^{R_c} \sqrt{2m[V(r)-E]}dr\right\}$$

S-factor

For the coulomb potential and assuming that $R_n \approx 0 \ll R_c$ the integral gives:

$$T_0 = e^{-2\pi\eta}, \quad \eta = \frac{Z_a Z_A e^2}{\hbar}\sqrt{\frac{m}{2E}}$$

η is the Sommerfeld parameter that accounts for tunneling through a coulomb barrier.

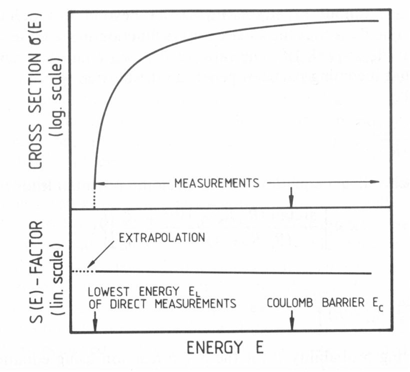
We can rewrite the cross section as:

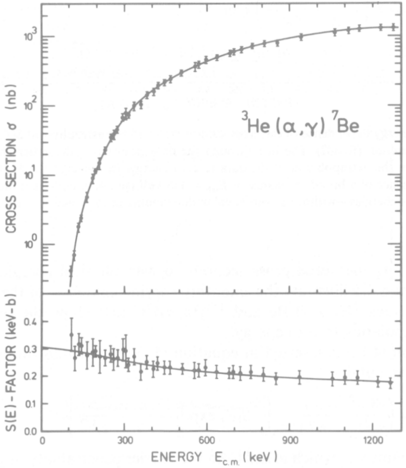
$$\sigma(E) = \frac{1}{E} S(E) e^{-2\pi\eta}$$

S is the so-called S-factor and accounts for the short distance dependence of the cross section on the nuclear potential. It is expected to be only mildly dependent on Energy.

S-factor

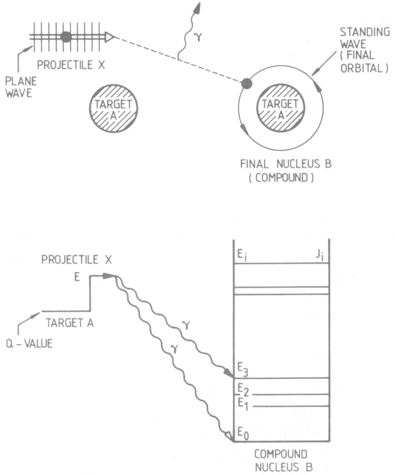
S factor makes possible accurate extrapolations to low energy.





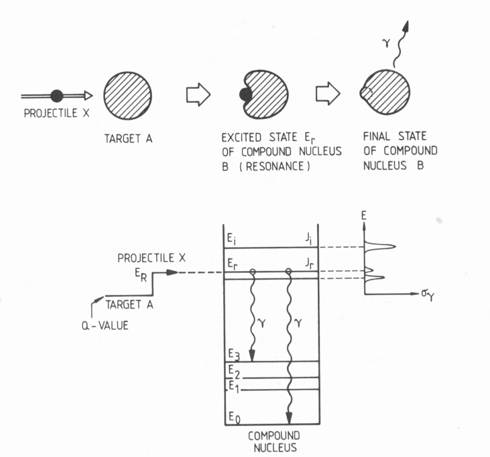
Direct reactions

So far we have discussed the so-called "direct reactions" in which the reaction proceeds directly to a bound nuclear state:



Resonant reactions

The cross section can also have contributions from resonances that can be seen like quasi-bound states. During the reaction a quasi-bound, compound, state forms that decays by particle or gamma emission.



Resonance cross section

The cross section for capture through an isolated resonance is given by the Breit-Wigner formula:

$$a + A \to C \to B + b$$

$$\sigma(E) = \pi \lambda\!\!\!^{-2} \frac{(2J_C + 1)(1 + \delta_{aA})}{(2J_a + 1)(2J_A + 1)} \frac{\Gamma_{aA}\Gamma_{bB}}{(E - E_r)^2 + (\Gamma/2)^2}, \qquad \lambda\!\!\!^{-} = \frac{1}{k} = \frac{\hbar}{p}$$

with $\Gamma = \Gamma_{aA} + \Gamma_{bB} + \ldots$ (sum over all partial widths for all possible decay channels). They depend on energy.

Particle width For charged particles is strongly dependent on energy due to tunneling through coulomb barrier.

Photon width Depends on the so called gamma strength function (will be discussed later)

Hauser-Feshbach cross section and reaction rate

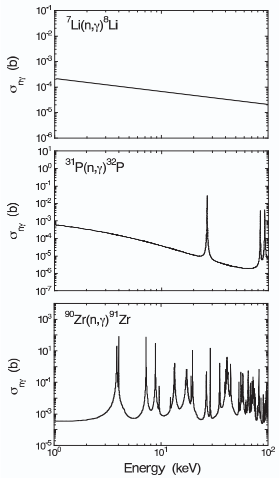


Fig. 3.30 Cross sections for neutron capture on ^{7}Li, ^{31}P, and ^{90}Zr versus energy. The curve in the upper panel shows a $1/v$ behavior, while resonances are visible in the middle and lower panels.

Christian Iliadis, *Nuclear Physics of Stars*

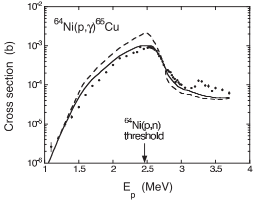


Fig. 2.29 Cross section versus bombarding energy for the ^{64}Ni(p,γ)^{65}Cu reaction. Beyond an energy of $\approx$ 2.5 MeV the endothermic ^{64}Ni(p,n)^{64}Cu reaction is energetically allowed. The sharp drop in the cross section at the neutron threshold reflects the decrease of the flux in all other decay channels of the compound nucleus ^{65}Cu. The curves show the results of Hauser–Feshbach statistical model calculations with (solid line) and without (dashed line) width fluctuation corrections. Reprinted from F. M. Mann et al., Phys. Lett. B, Vol. 58, p. 420 (1975). Copyright (1975), with permission from Elsevier.

- With increasing mass number reactions are determined by a larger number of resonances
- Often it is not possible to experimentally resolve resonances. Astrophysical reaction rate is an energy average over many resonances.
- Hauser-Feshbach provides and statistically averaged cross section from the contribution of many resonances in an energy interval.

Hauser-Feshbach cross section

The Hauser-Feshbach expression for the cross section of an (n, γ) reaction proceeding from the target nucleus i in the state μ with spin J_i^μ and parity π_i^μ to a final state ν with spin J_m^ν and parity π_m^ν in the residual nucleus m via a compound state with excitation energy E, spin J, and parity π is given by

$$\sigma_{(n,\gamma)}^{\mu\nu}(E_{i,n}) = \frac{\pi\hbar^2}{2M_{i,n}E_{i,n}} \frac{1}{(2J_i^\mu + 1)(2J_n + 1)} \sum_{J,\pi} (2J + 1) \frac{T_n^\mu T_\gamma^\nu}{T_n^\mu + T_\gamma^\nu}$$

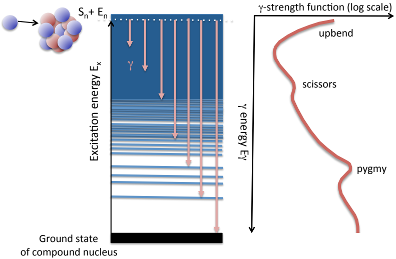
where $E_{i,n}$ and $M_{i,n}$ are the center-of-mass energy and the reduced mass for the initial system. $J_n = 1/2$ is the neutron spin. Normally we have situations in which $T_n \gg T_\gamma$. The transmission coefficients, multipolarity XL, are related to the average decay width and level density (ρ) or the gamma strength function, f_{XL}:

$$T_{XL} = 2\pi\rho\langle\Gamma_{XL}\rangle = 2\pi E_\gamma^{2L+1} f_{XL}$$

Structure dipole gamma-strength function

The gamma-decay is dominated by dipole transitions. Total transmission coefficient (sum final bound states)

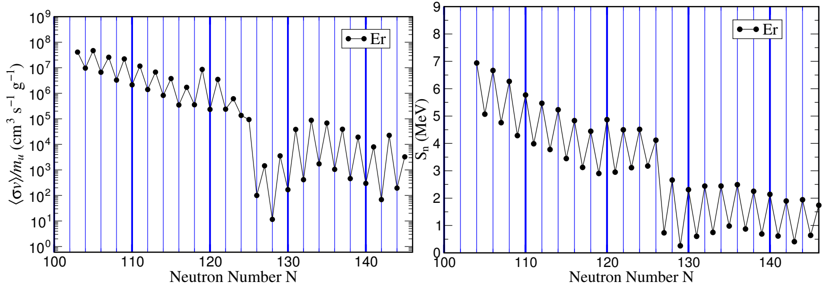
$$T_1 = \sum_{\nu} 2\pi E_{\gamma}^3 f_{1,\nu} = 2\pi \int E_{\gamma}^3 f_1(E_{\gamma}) \rho(E_{\gamma}) dE_{\gamma}$$



Favored decay energy determined by a competition between level density, ρ, and gamma-strength function.

Systematics $\langle\sigma v\rangle$ and neutron separation energies

Neutron capture rates reflect the behavior of neutron separation energies



Inverse reactions

Let's have the reaction

$$a + A \rightarrow B + \gamma \quad Q = m_a + m_A - m_B$$

We are interested in the inverse reaction. One can use detailed-balance to determine the inverse rate. Simpler using the concept of chemical equilibrium.

$$\frac{dn_a}{dt} = -n_a n_A \langle \sigma v \rangle_{aA} + (1 + \delta_{aA}) n_B \lambda_\gamma = 0$$

$$\left(\frac{n_a n_A}{n_B} \right)_{\text{eq}} = (1 + \delta_{aA}) \frac{\lambda_\gamma}{\langle \sigma v \rangle_{aA}}$$

Using equilibrium condition for chemical potentials: $\mu_a + \mu_A = \mu_B$

$$\mu(Z, A) = m(Z, A)c^2 + kT \ln \left[\frac{n(Z, A)}{G_{(Z,A)}(T)} \left(\frac{2\pi\hbar^2}{m(Z, A)k_B T} \right)^{3/2} \right], \quad G_{(Z,A)}(T) = \sum_i (2J_i + 1) e^{-E_i/(kT)}$$

Inverse reactions

One obtains:

$$\left(\frac{n_a n_A}{n_B}\right)_{\text{eq}} = \frac{G_a G_A}{G_B}\left(\frac{m_a m_A}{m_B}\right)^{3/2}\left(\frac{k_B T}{2\pi\hbar^2}\right)^{3/2} e^{-Q/k_B T}$$

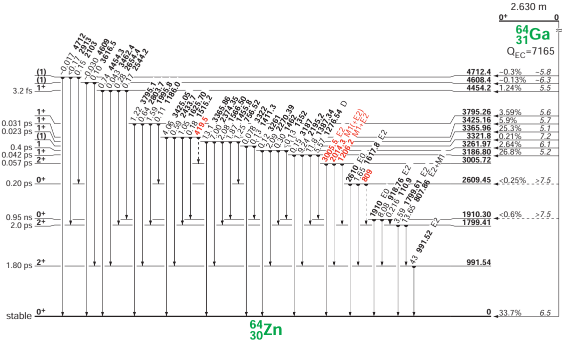
Finally, we obtain:

$$\lambda_\gamma = \frac{G_a G_A}{(1+\delta_{aA})G_B}\left(\frac{m_a m_A}{m_B}\right)^{3/2}\left(\frac{k_B T}{2\pi\hbar^2}\right)^{3/2} e^{-Q/k_B T}\langle\sigma v\rangle$$

For a reaction $a + A \to B + b$ ($Q = m_a + m_A - m_B - m_b$):

$$\langle\sigma v\rangle_{bB} = \frac{(1+\delta_{bB})}{(1+\delta_{aA})}\frac{G_a G_A}{G_b G_B}\left(\frac{m_a m_A}{m_b m_B}\right)^{3/2} e^{-Q/k_B T}\langle\sigma v\rangle_{aA}$$

Beta-decay

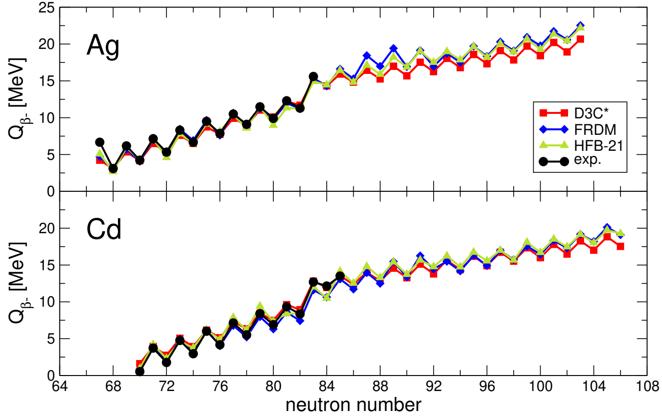


$$\lambda_\beta = \frac{\ln 2}{K} f(Q)[B(F) + B(GT)], \qquad K = \frac{2\pi^3 (\ln 2)\hbar^7}{G_F^2 V_{ud}^2 g_V^2 m_e^5 c^4} = 6147.0 \pm 2.4\ s$$

- $f(Q)$ phase space function ($\sim Q^5$).
- $B(F)$ Fermi matrix element ($\sim \langle f | \sum_k t_+^k | i \rangle$).
- $B(GT)$ Gamow-Teller matrix element ($\sim \langle f | \sum_k \boldsymbol{\sigma} t_+^k | i \rangle$).

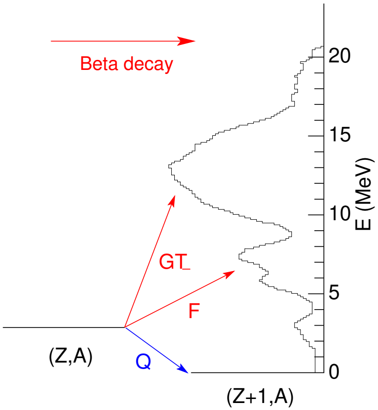
Systematics beta-decay Q-values

The decay Q-value increases with neutron excess.



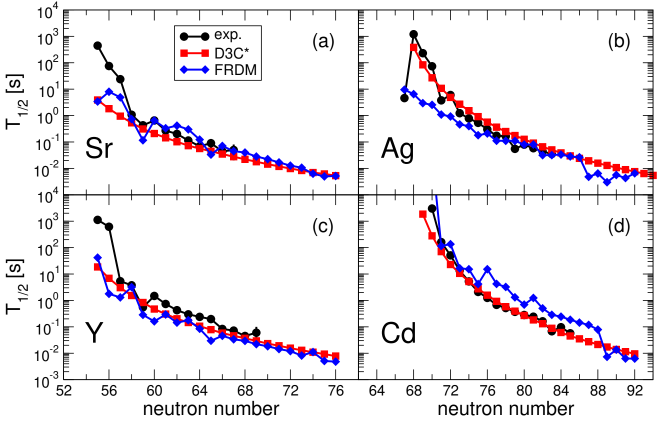
Structure Gamow-Teller strength

- Gamow-Teller strength is characterized by the presence of a resonance at excitation energies around 10-15 MeV.
- With increasing neutron-excess a larger fraction of strength is in the decay window.
- Low-lying strength is rather sensitive to correlations.



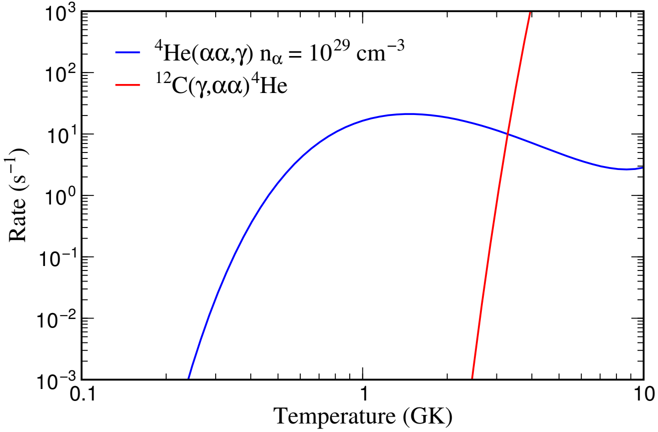
Systematics half-lives

Half-lives decrease with neutron excess.

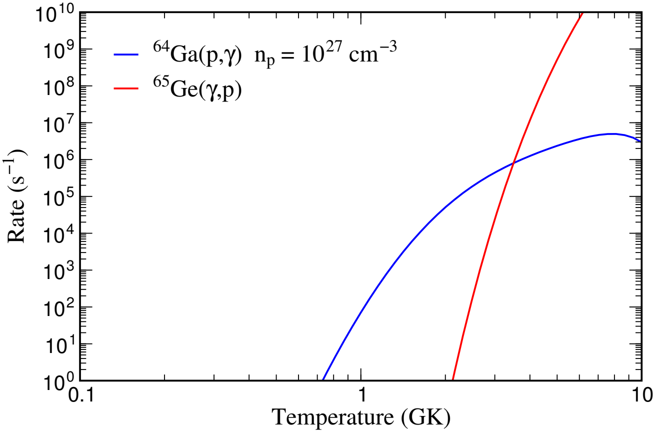


T. Marketin, L. Huther, and G. Martínez-Pinedo, Phys. Rev. C **93**, 025805 (2016).

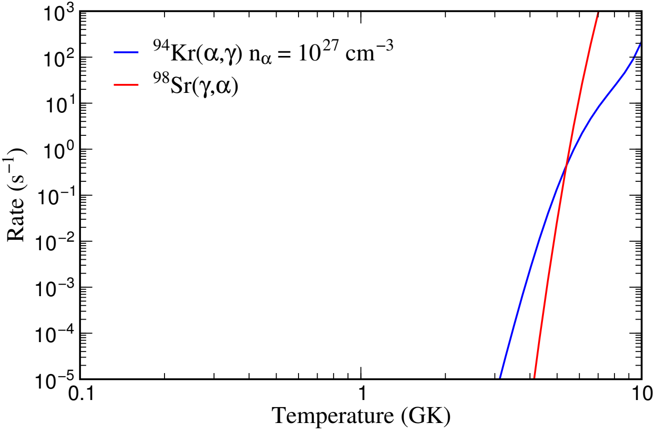
Rate Examples: $^4\text{He}(\alpha\alpha,\gamma)$



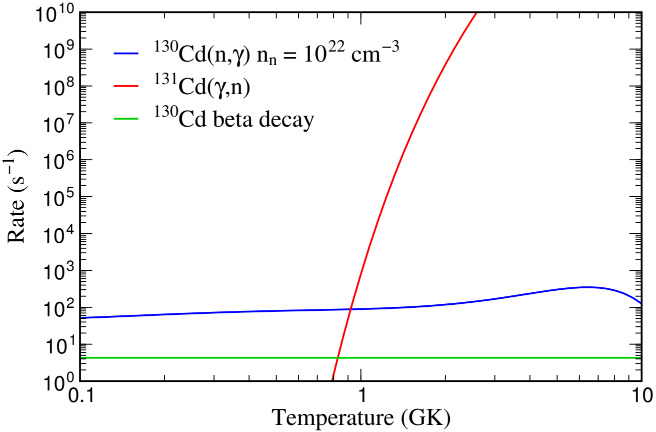
Rate Examples: (p, γ)



Rate Examples: (α, γ)



Rate examples: (n, γ)



R-process sites

- Any r-process site should be able to produce both the "seed" nuclei where neutrons are captured and the neutrons that drive the r-process. The main parameter describing the feasibility of a site to produce r-process nuclei is the neutron-to-seed ratio: n_n/n_{seed}.
- If the seed nuclei have mass number A_{seed} and we have n_n/n_{seed} neutrons per seed, the final mass number of the nuclei produced will be $A = A_{\text{seed}} + n_n/n_{\text{seed}}$.
- For example, taking $A_{\text{seed}} = 90$ we need $n_n/n_{\text{seed}} = 100$ if we want to produce the 3rd r-process peak ($A \sim 195$) and $n_n/n_{\text{seed}} = 150$ to produce U and Th.

R-process sites

In an astrophysical site there are only two possible ways to achieve large neutron-to-seeds:

1. Let us consider high temperature neutron-rich matter with high entropy that it is ejected at high velocities. As the material expands α particles will be formed. However, the build up of heavy nuclei by 3-body reactions becomes very unefficient by two reasons: 1) Too many photons per nucleon due to the high entropy, 2) Too litle time to produce heavy nuclei due to the fast expansion. It means that we will have an α-rich freeze out with a few heavy nuclei produced and many neutrons left ($Y_\alpha \approx Y_e/2$, $Y_n \approx 1 - 2Y_e$). This is commonly denoted as "high entropy" r-process
2. Let us consider matter very high density matter with low entropies. Due to the high densities electrons have large fermi energies and will drive the composition very neutron rich. At some point the neutron drip line is reached and nuclei start to "drip" neutrons. This is the situation in the crust of neutron stars where densities are 10^{12-13} g cm^{-3} and $Y_e \sim 0.05$: $Y_n = 1 - \langle A \rangle Y_e/\langle Z \rangle$, $Y_s = Y_e/\langle Z \rangle$; $Y_n/Y_s = \langle Z \rangle/Y_e - \langle A \rangle$; $Y_n/Y_s \sim 500 - 2000$. This is commonly denoted as "low entropy" r-process.

r-process nucleosynthesis relevant parameters

Independently of the astrophysical site the nucleosynthesis is sensitive to a few parameters that determine the neutron-to-seed ratio and the heavier elements that can be produced:

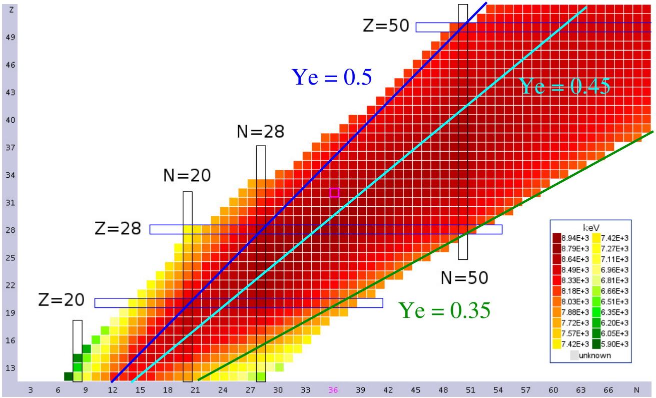
$$A_f = A_i + n_s, \quad n_s = n_n/n_{\text{seed}} \sim s^3/(Y_e^3 \tau_{\text{dyn}})$$

Y_e The lower the value of Y_e more neutrons are available and the larger n_s

entropy Large entropy $s \sim T^3/\rho$, means low density and high temperature (large amount of photons). Both are detrimental to the build up of seeds by 3-body reactions.

expansiton time scale The faster the matter expands, smaller τ_{dyn}, the less time one has to build up seeds

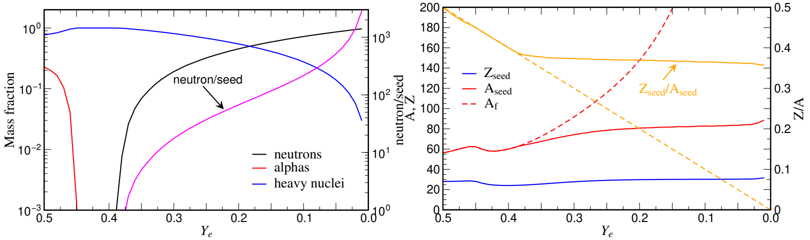
Dependence on Y_e



With reduced Y_e the peak of the nuclear abundance distribution moves from ^{56}Ni ($Y_e \sim 90$) to heavier neutron rich nuclei (^{90}Zr for $Y_e \sim 0.45$). For low Y_e becomes energetically favourable to have free neutrons.

Dependence abundances (at 3 GK) on Y_e

For moderate entropies r process depends mainly on Y_e.



For neutron-rich moderate entropy ejecta we have:

$$n_s = A\left(\frac{Z}{A}\frac{1}{Y_e} - 1\right)$$

Calculation by Bowen Jiang

r process calculations

r process calculations require to solve the system of differential equations:

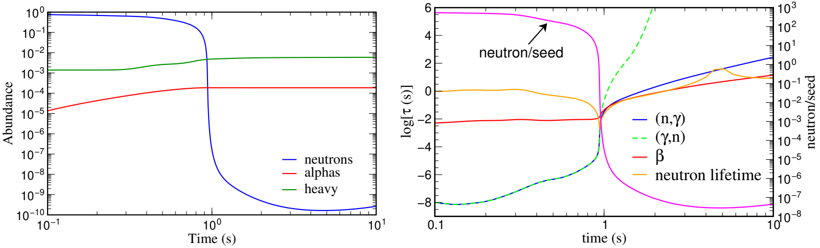
$$
\begin{aligned}
\frac{dY(Z,A)}{dt} = & \left(\frac{\rho}{m_u}\right)\langle\sigma v\rangle_{Z,A-1} Y_n Y(Z,A-1) + \lambda_\gamma(Z,A+1)Y(Z,A+1) \\
& + \sum_{j=0}^{J} \lambda_{\beta jn}(Z-1,A+j)Y(Z-1,A+j) \\
& - \left(\left(\frac{\rho}{m_u}\right)\langle\sigma v\rangle_{Z,A} Y_n + \lambda_\gamma(Z,A) + \sum_{j=0}^{J} \lambda_{\beta jn}(Z,A)\right) Y(Z,A)
\end{aligned}
$$

$$
\begin{aligned}
\frac{dY_n}{dt} = & - \sum_{Z,A} \left(\frac{\rho}{m_u}\right)\langle\sigma v\rangle_{Z,A} Y_n Y(Z,A) \\
& + \sum_{Z,A} \lambda_\gamma(Z,A)Y(Z,A) \\
& + \sum_{Z,A} \left(\sum_{j=1}^{J} j\lambda_{\beta jn}(Z,A)\right) Y(Z,A)
\end{aligned}
$$

We are neglecting fission

Evolution during r process

Example of r process calculation for very neutron rich ejecta (Based on trajectory from merger simulation from A. Bauswein)



$$\frac{1}{\tau_n} = \frac{1}{n_s}\left(\frac{1}{\tau_{(n,\gamma)}} - \frac{1}{\tau_{(\gamma,n)}}\right)$$

The last phase of the r process when the neutron-to-seed decreases very fast is known as freeze-out. During this phase neutron captures and beta-decays compete with each other. It is during this phase that the final abundances are shaped.

From Meng-Ru Wu

$(n,\gamma) \rightleftarrows (\gamma,n)$ equilibrium

Neutron capture reactions proceed much faster than beta-decays and an $(n,\gamma) \rightleftarrows (\gamma,n)$ equilibrium is achieved

$$\mu(Z,A+1) = \mu(Z,A) + \mu_n$$

$$\frac{Y(Z,A+1)}{Y(Z,A)} = n_n \left(\frac{2\pi\hbar^2}{m_u k_B T}\right)^{3/2} \left(\frac{A+1}{A}\right)^{3/2} \frac{G(Z,A+1)}{2G(Z,A)} \exp\left[\frac{S_n(Z,A+1)}{k_B T}\right]$$

Only even-even nuclei participate in the path so we can write:

$$\frac{Y(Z,A+2)}{Y(Z,A)} = n_n^2 \left(\frac{2\pi\hbar^2}{m_u k_B T}\right)^{3} \left(\frac{A+2}{A}\right)^{3/2} \frac{G(Z,A+2)}{4G(Z,A)} \exp\left[\frac{S_{2n}(Z,A+2)}{k_B T}\right]$$

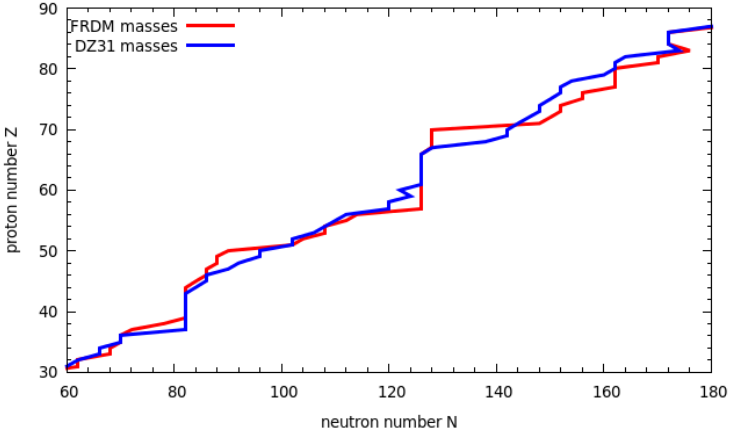
The maximum of the abundance defines the r-process path:

$$S_{2n}^0(\text{MeV}) = \frac{2T_9}{5.04}\left(34.075 - \log n_n + \frac{3}{2}\log T_9\right)$$

For $n_n = 5\times 10^{21}\ \text{cm}^{-3}$ and $T = 1.3$ GK corresponds at $S_{2n}^0 = 6.46$ MeV, $S_n^0 = S_{2n}^0/2 = 3.23$ MeV,

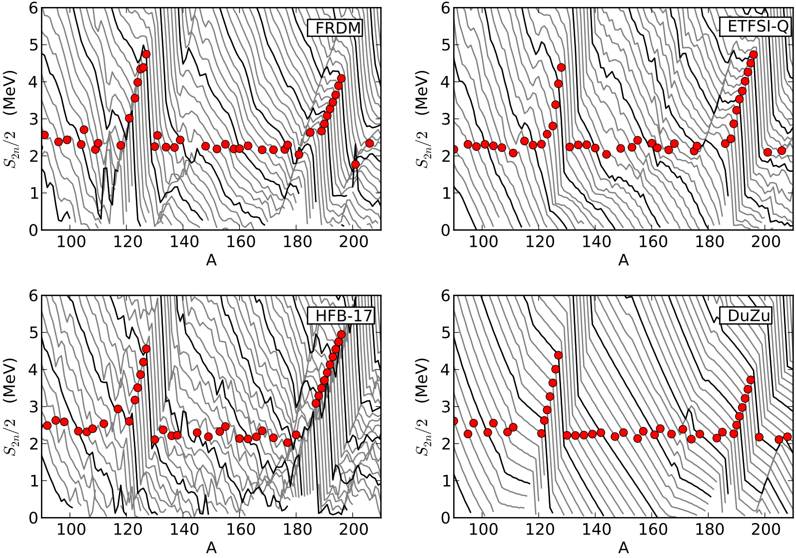
r process path

Path depends on nuclear masses



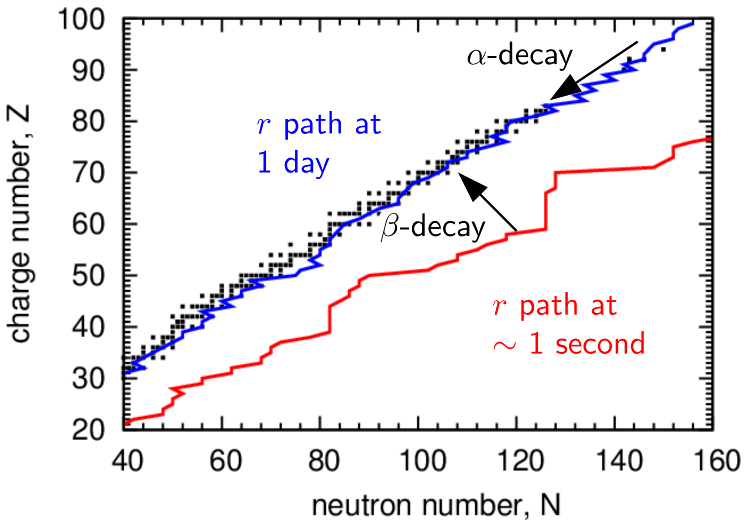
Two-neutron separation energies

r-process path defines by those nuclei with: $S_{2n}(Z, A) \gtrsim S_{2n}^0$



A. Arcones, GMP, Phys. Rev. C **83**, 045809 (2011)

path dependence on time



Beta-flow equilibrium

Assuming $(n,\gamma) \rightleftarrows (\gamma,n)$ equilibrium, it is sufficient to compute the time evolution of the total abundance for an isotopic chain

$$Y(Z) = \sum_A Y(Z,A)$$

The differential equation reduces to

$$\frac{dY(Z)}{dt} = \lambda_\beta(Z-1)Y(Z-1) - \lambda_\beta(Z)Y(Z)$$

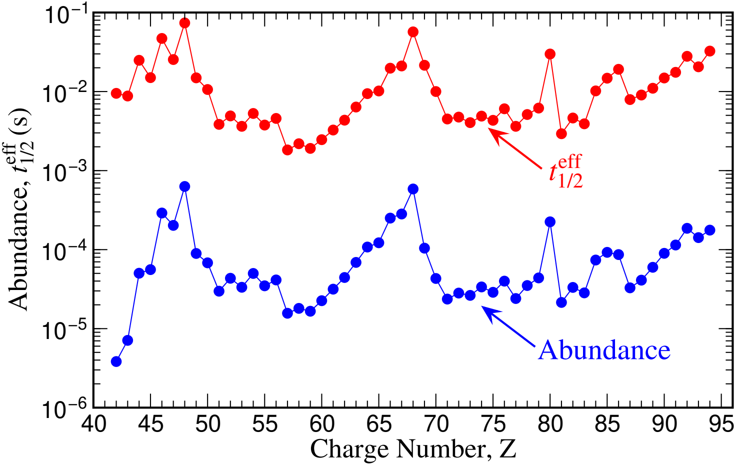
with

$$\lambda_\beta(Z) = \frac{1}{Y(Z)} \sum_A \lambda_\beta(Z,A)Y(Z,A)$$

Only beta-decays are necessary to determine the evolution. If the duration of the r process is larger than the beta-decay lifetimes an equilibrium is reaches denotes as steady β flow equilibrium that satisfies for each Z value

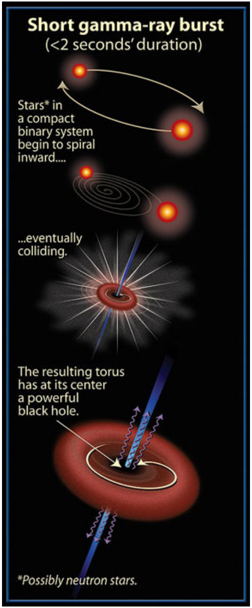
$$\lambda_\beta(Z-1)Y(Z-1) = \lambda_\beta(Z)Y(Z), \quad \text{i.e. } Y(Z) \propto \tau_\beta(Z)$$

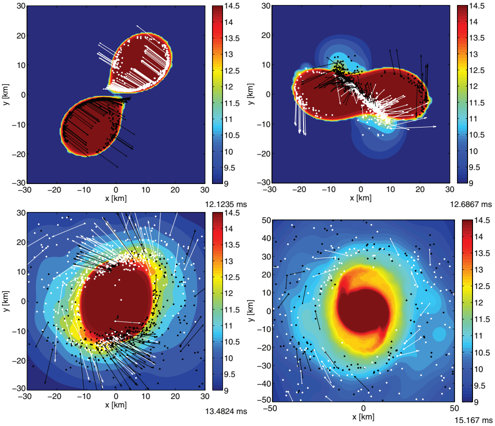
Abundances vs beta lifetimes



$$\frac{1}{t_{1/2}^{\mathrm{eff}}(Z)} = \frac{1}{\sum_N Y(Z,N)} \sum_N \frac{Y(Z,N)}{t_{1/2}(Z,N)}$$

Neutron star mergers: Short gamma-ray bursts and r-process



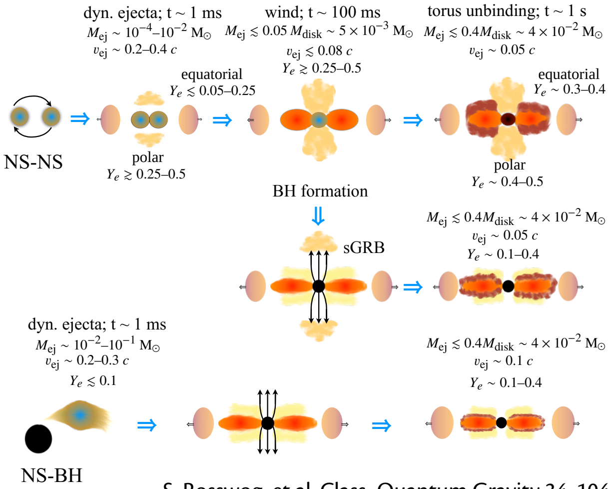


Basuswein, Goriely, Janka, ApJ 773, 78 (2013)

- Mergers are associated with short-gamma ray bursts.
- They are also promising sources of gravitational waves.
- Observational signatures of the r-process?

Merger channels and ejection mechanism

In mergers we deal with a variety of initial configurations (netron-star neutron-star vs neutron-star black-hole) with additional variations in the mass-ratio. The evolution after the merger also allows for further variations.

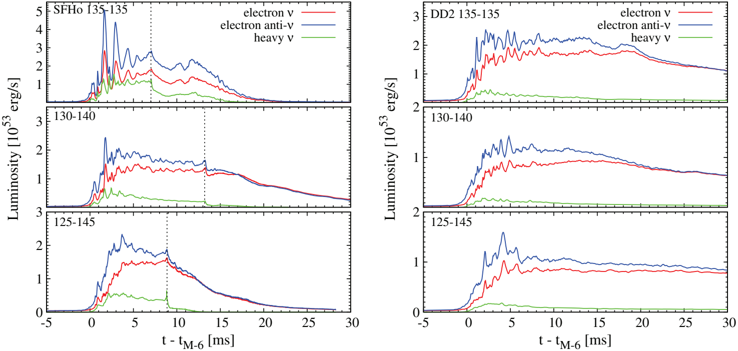


S. Rosswog, et al, Class. Quantum Gravity 34, 104001 (2017).

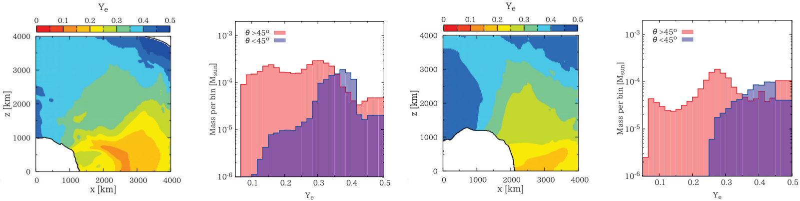
Impact of neutrinos on NS-NS mergers

Sekiguchi+, PRD **93**, 124046 (2016)

Mergers are characterized by large $\bar{\nu}_e$ luminosities



Shibata+, PRD **96**, 123012 (2017)



Why is Y_e modified?

Y_e is mainly determined by the competition between

$$\nu_e + n \rightleftarrows p + e^-$$
$$\bar{\nu}_e + p \rightleftarrows n + e^+$$

The evolution of Y_e is given by:

$$\frac{dY_e}{dt} = \lambda_{\nu_e n} Y_n - \lambda_{\bar{\nu}_e p} Y_p = \lambda_{\nu_e n} - (\lambda_{\nu_e n} + \lambda_{\bar{\nu}_e p}) Y_e,$$

using $Y_e = Y_p$ and $Y_n = 1 - Y_n$. If matter is exposed long time enough to neutrino fluxes it will reach an equilibrium given by:

$$Y_e^{\mathrm{eq}} = \frac{\lambda_{\nu_e n}}{\lambda_{\nu_e n} + \lambda_{\bar{\nu}_e p}}.$$

We need to estimate the neutrino absorption rates.

neutrino absorption rates on nucleons

If we assume that the produced electron (positron) is extremelly relativistic ($E_e = p_e c$) the cross section for (anti)neutrino absorption is:

$$\sigma(E_\nu) = \sigma_0 (E_\nu \pm \Delta)^2$$

plus sign for neutrinos and minus sign for antineutrinos, $\Delta = m_n c^2 - m_p c^2 = 1.2933\ \mathrm{MeV}$, and

$$\sigma_0 = \frac{(1+3g_A^2) G_F^2 V_{ud}^2}{\pi \hbar^4 c^4} = 9.33(4) \times 10^{-44}\ \mathrm{cm^2\ MeV^{-2}}$$

For the absorption rate we need to integrate over the neutrino spectrum and multiply by the neutrino flux ($L_\nu/(4\pi r^2 \langle E_\nu \rangle)$) obtaining:

$$\lambda_{\nu_e n} = \frac{L_{\nu_e}}{4\pi r^2} \sigma_0 \left(\varepsilon_{\nu_e} + 2\Delta + \frac{\Delta^2}{\langle E_{\nu_e} \rangle} \right)$$

$$\lambda_{\bar{\nu}_e p} = \frac{L_{\bar{\nu}_e}}{4\pi r^2} \sigma_0 \left(\varepsilon_{\bar{\nu}_e} - 2\Delta + \frac{\Delta^2}{\langle E_{\bar{\nu}_e} \rangle} \right)$$

with $\varepsilon_\nu = \langle E_\nu^2 \rangle / \langle E_\nu \rangle$.

Equilibrium Y_e

The equilibrium Y_e^{eq} can be expressed as:

$$Y_e^{\text{eq}} = \left[1 + \frac{L_{\bar{\nu}_e}}{L_{\nu_e}} \frac{\varepsilon_{\bar{\nu}_e} - 2\Delta + \Delta^2/\langle E_{\bar{\nu}_e}\rangle}{\varepsilon_{\nu_e} + 2\Delta + \Delta^2/\langle E_{\nu_e}\rangle}\right]^{-1}$$

In order to achieve $Y_e < 0.5$ we need:

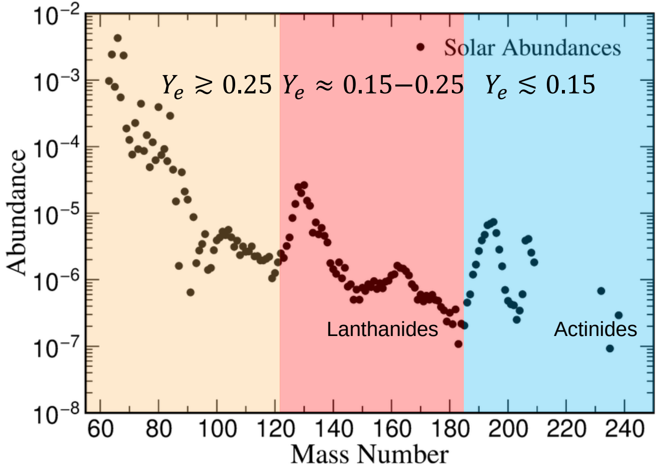
$$Y_e^{\text{eq}} = \frac{\lambda_{\nu_e n}}{\lambda_{\nu_e n} + \lambda_{\bar{\nu}_e p}} < 1/2 \Rightarrow \lambda_{\bar{\nu}_e p} > \lambda_{\nu_e n}$$

This can be translated to a condition on the average energies:

$$\varepsilon_{\bar{\nu}_e} - \varepsilon_{\nu_e} > 4\Delta - \left[\frac{L_{\bar{\nu}_e}}{L_{\nu_e}} - 1\right][\varepsilon_{\bar{\nu}_e} - 2\Delta]$$

Dependence nucleosynthesis on Y_e

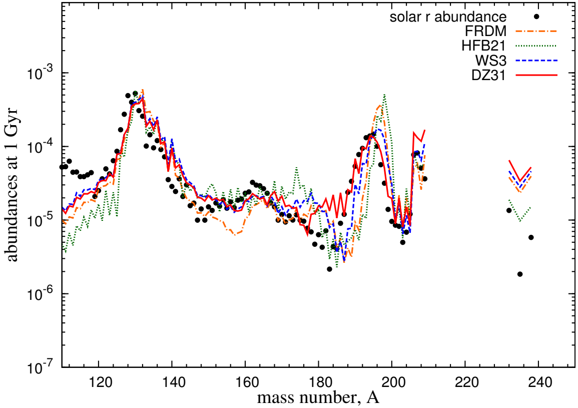
Nucleosynthesis depends on neutron richness of ejecta



The relevant nuclear physics depends on the particular conditions.

Nucleosynthesis low Y_e ejecta

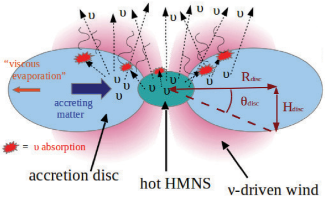
Mendoza-Temis, Wu, Langanke, GMP, Bauswein, Janka, PRC **92**, 055805 (2015)



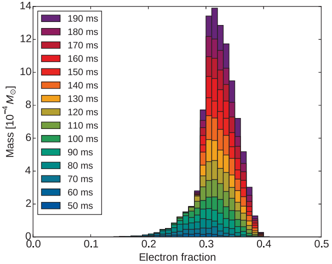
- Robustness astrophysical conditions, strong sensitivity to nuclear physics
- Second peak ($A \sim 130$) sensitive to fission yields.
- Third peak ($A \sim 195$) sensitive to masses (neutron captures) and beta-decay half-lives.

Post-merger Nucleosynthesis (NS remnant)

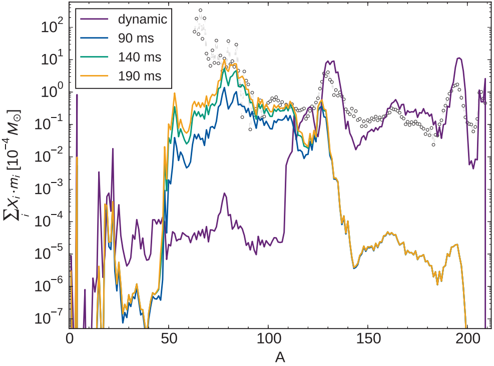
An Hypermassive neutron star produces large neutrino fluxes that drive the composition to moderate neutron rich ejecta.



Perego, *et al*, MNRAS 443, 3134 (2014)



Martin, *et al*, ApJ 813, 2 (2015)

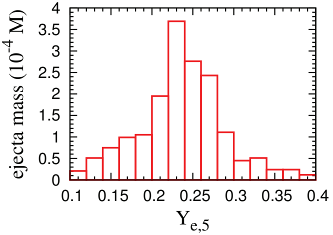


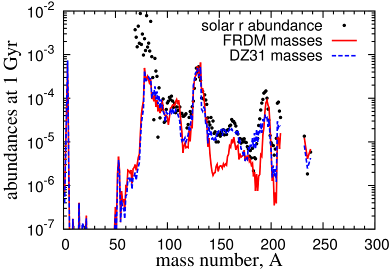
Only nuclei with $A < 120$ are produced (no lanthanides, blue kilonova).

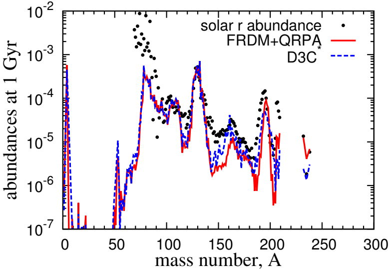
See also Lippuner et al, MNRAS **472**, 904 (2017)

Post-merger nucleosynthesis (BH remnant)

- Accretion disk around BH ejects relatively neutron-rich matter [Fernández, Metzger, MNRAS 435, 502 (2013)]
- Produces all r-process nuclides (Lanthanide/Actinide rich ejecta) [Wu *et al*, MNRAS 463, 2323 (2016)]



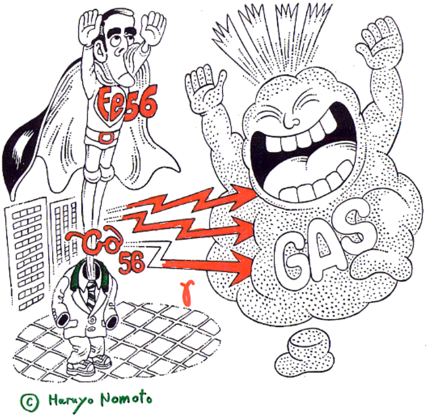


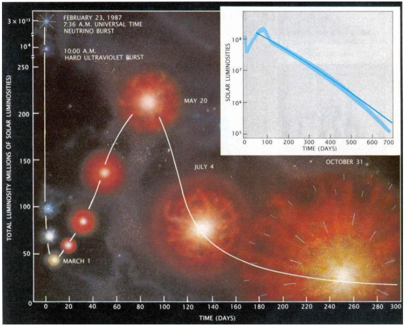


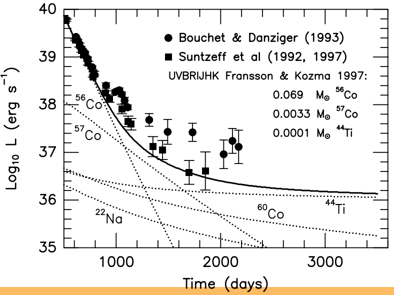
See also Just et al, MNRAS 448, 541 (2015), Siegel and Metzger PRL 119, 231102 (2017)

Electromagnetic signatures of nucleosynthesis

Supernova light curves follow the decay of ^{56}Ni ($t_{1/2} = 6$ d) and later ^{56}Co ($t_{1/2} = 77$ d)

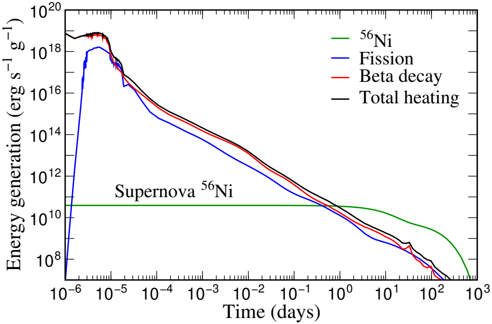






Energy production from r-process ejecta

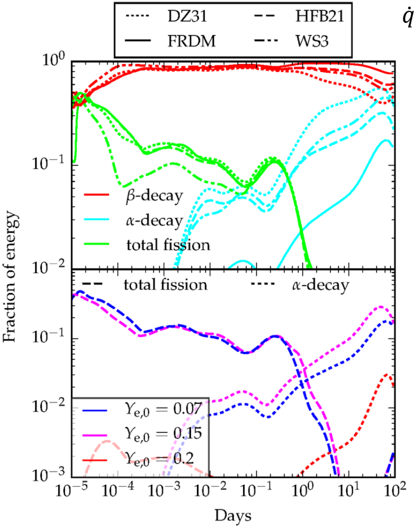
The decay of r process products produces energy following a power law $\varepsilon \sim t^{-1.3}$ (Way & Wigner 1948, Metzger *et al.* 2010). Many nuclei decaying at the same time heating up the ejecta



We expect an electromagnetic transient (Li & Paczyńsky 1998) whose properties depend on:

- Energy production rate
- Efficiency energy is absorbed by gas (thermalization efficiency)
- Opacity gas (depends on composition, presence of lanthanides/actinides)

Energy production and thermalization

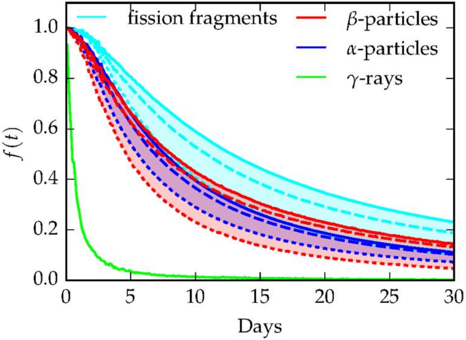


$$\dot{q}(t) = \sum_k f_k(t)\dot{\varepsilon}_k(t)$$

$\dot{\varepsilon}_k(t)$ energy emitted in particle k

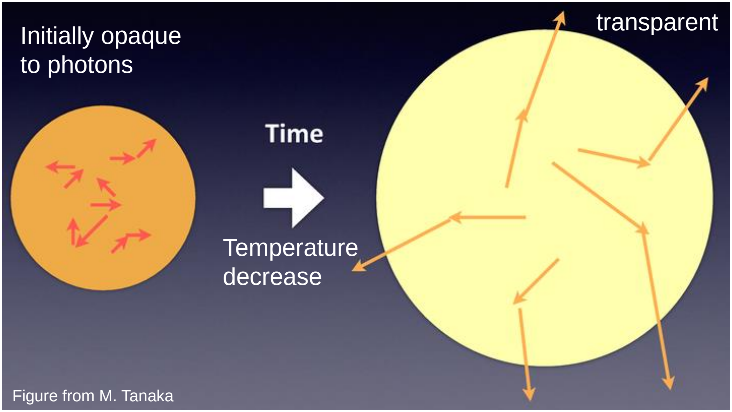
$f_k(t)$ thermalization efficiency particle k

Thermalization depends on particle, ejecta dynamics, magnetic field, …



See Barnes, Kasen, Wu, GMP, ApJ 829, 110 (2016); Kasen & Barnes, arXiv:1807.03319

Impact of opacity

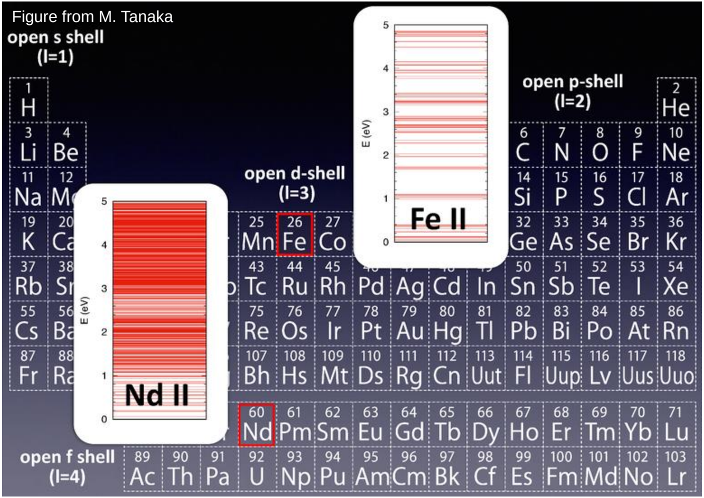


The transition from an opaque to transparent regime depends on the interaction probability of the photons (opacity). Depends on the structure of the atoms.

Low opacity: early emission from hot material at short wavelengths (blue)

High opacity: late emission from colder material at longer wavelengths (red)

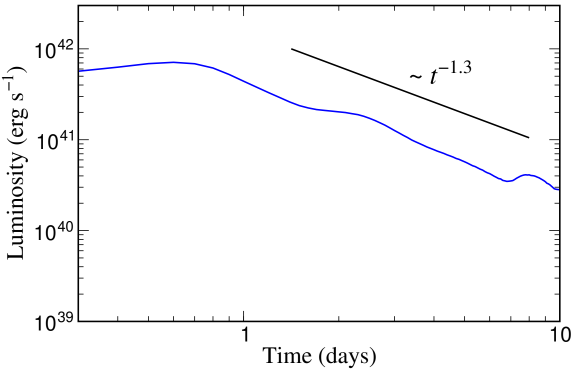
Impact Lanthanides



Large number of states of Lanthanides/Actinides leads to a high opacity

Barnes & D. Kasen, Astrophys. J. 775, 18 (2013);Tanaka & Hotokezaka, Astrophys. J. 775, 113 (2013).

Kilonova: Electromagnetic signature of the r process



Luminosity equivalent to 1000 novas (kilonova) in timescales of days.
Depends on amount of ejected material, velocity and composition (opacity)

Metzger, GMP, Darbha, Quataert, Arcones *et al*, MNRAS **406**, 2650 (2010)

Simple Kilonova model

Light curve is expected to peak when photon diffusion time is comparable to elapsed time (Metzger et al 2010, Kasen et al 2017)

$$t_{\mathrm{diff}} = \frac{\rho \kappa R^2}{3\, c}, \qquad \rho = \frac{M}{4\pi R^3/3}, \qquad R = vt$$

$$t_{\mathrm{peak}} \approx \left(\frac{\kappa M}{4\pi c v}\right)^{\frac{1}{2}} \approx 1.5\ \mathrm{days} \left(\frac{M}{0.01\ M_{\odot}}\right)^{\frac{1}{2}} \left(\frac{v}{0.01c}\right)^{-\frac{1}{2}} \left(\frac{\kappa}{1\ \mathrm{cm^2\ g^{-1}}}\right)^{\frac{1}{2}}$$

The Luminosity is $L(t) \approx M\, \dot{\varepsilon}(t)$, $\dot{\varepsilon}(t) \approx 10^{10} \left(\frac{t}{1\ \mathrm{day}}\right)^{-\alpha} \mathrm{erg\ s^{-1}\ g^{-1}}$

$$L_{\mathrm{peak}} \approx 1.1 \times 10^{41} \mathrm{erg\ s^{-1}} \left(\frac{M}{0.01\ M_{\odot}}\right)^{1-\frac{\alpha}{2}} \left(\frac{v}{0.01c}\right)^{\frac{\alpha}{2}} \left(\frac{\kappa}{1\ \mathrm{cm^2\ g^{-1}}}\right)^{-\frac{\alpha}{2}}$$

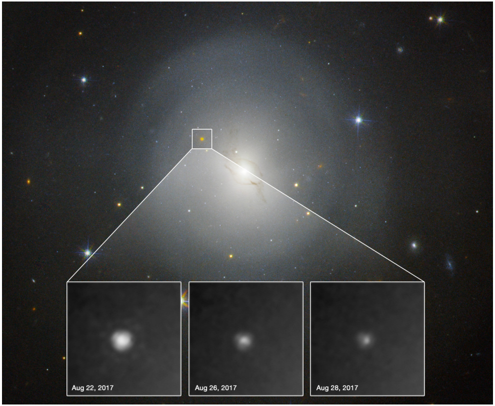
Very sensitive to atomic opacity

$\kappa \approx 1\ \mathrm{cm^2\ g^{-1}}$, light r process material (blue emission)

$\kappa \approx 10\ \mathrm{cm^2\ g^{-1}}$, heavy (lanthanide/actinide rich) r process (red emis.)

AT 2017 gfo: electromagnetic signature from r process

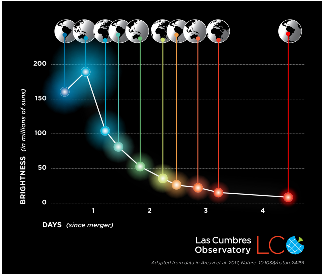
In-situ signature of r process nucleosynthesis

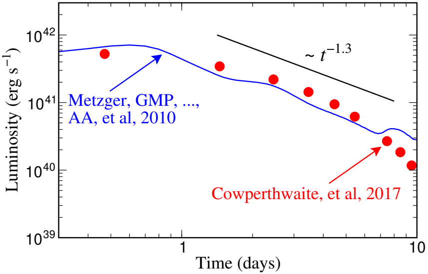


NASA and ESA. N. Tanvir (U. Leicester), A. Levan (U. Warwick), and A. Fruchter and O. Fox (STScI)

- Novel fastly evolving transitent
- Signature of statistical decay of fresly synthesized r process nuclei

AT 2017 gfo: interpretation

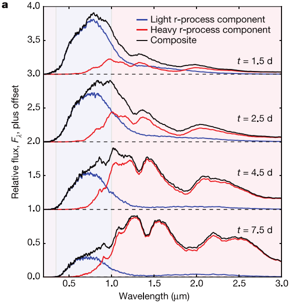




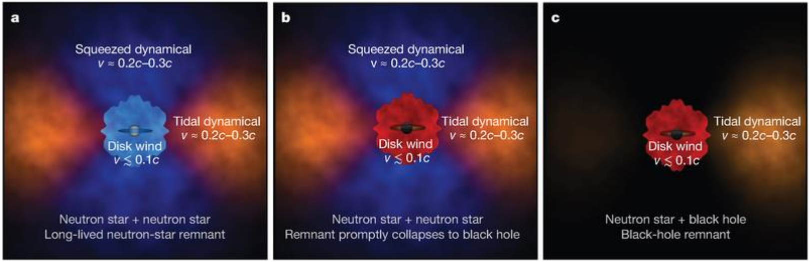
- Time evolution determined by the radioactive decay of r-process nuclei
- Two components:
  - blue dominated by light elements ($Z < 50$)
  - Red due to presence of Lanthanides ($Z = 57$–71) and/or Actinides ($Z = 89$–103)
- Likely source of heavy elements including Gold, Platinum and Uranium

Two components model

Kasen et al, Nature 551, 80 (2017)

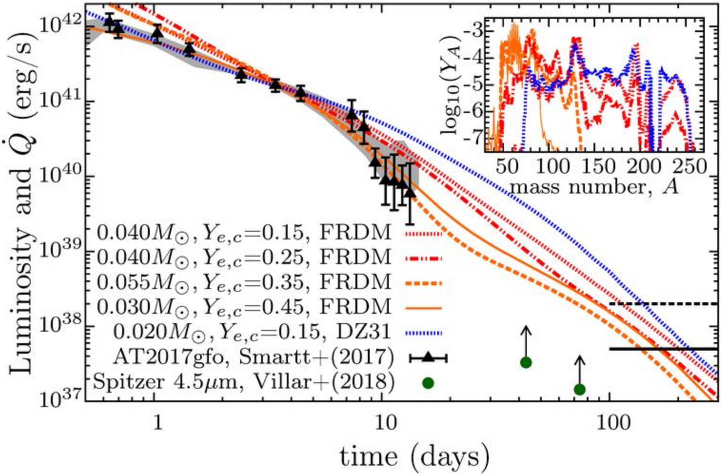


- Blue component from polar ejecta subject to strong neutrino fluxes (light r process)
  $M = 0.025\, M_\odot, v = 0.3c, X_{\text{lan}} = 10^{-4}$
- Red component disk ejecta after NS collapse to a black hole (includes both light and heavy r process)
  $M = 0.04\, M_\odot, v = 0.15c, X_{\text{lan}} = 10^{-1.5}$



Nuclear fingerprints light curve

Can we identify particular nuclear signatures in the light curve?



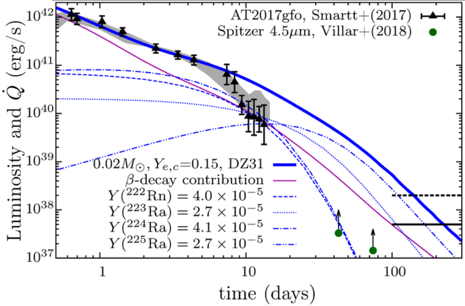
Observations between 10 and 100 days are sensitive to composition.
Light curve becomes dominated by individual decays

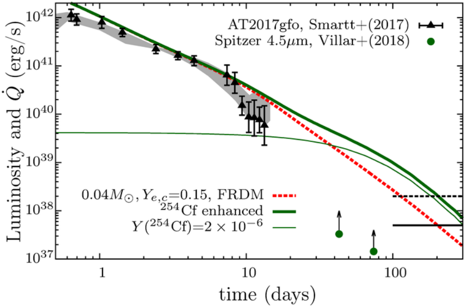
Wu, Barnes, GMP, Metzger, PRL 122, 062701 (2019)

Signature dominating decay chains

Isotope	Decay channel	$t_{1/2}$ (d)	Q (MeV)	E_α (MeV)	E_e (MeV)	E_γ (MeV)
^{224}Ra	$\alpha\beta^-$ to ^{208}Pb	3.6319(23)	30.875	26.542	0.891	1.474
^{222}Rn	$\alpha\beta^-$ to ^{210}Pb	3.8215(2)	23.826	19.177	0.949	1.715
^{225}Ra	β^-	14.9(2)	0.356	-	0.097	0.012
^{225}Ac	$\alpha\beta^-$ to ^{209}Bi	10.0(1)	30.196	27.469	0.632	0.046
^{223}Ra	$\alpha\beta^-$ to ^{207}Pb	11.43(5)	29.986	26.354	0.937	0.304

Isotope	Decay channel	$t_{1/2}$ (d)	Q (MeV)	$E_{\rm Kinetic}$ (MeV)	E_n (MeV)	E_γ (MeV)
^{254}Cf	Fission	60.5(2)	-	185(2)	-	-





Decline observed light curve at 10 days suggest an upper limit of 0.01 $M_\odot$ of U and Th

Wu, Barnes, GMP, Metzger, , PRL 122, 062701 (2019)

Summary

- Heavy elements are observed at very early times in Galactic history. Produced by a primary process that creates both neutrons and seeds.
- Neutron star mergers are likely the site where the “main r process” takes place.
- Radioactive decay of r-process ejecta produces an electromagnetic transient observed for the first time after GW170817.
- Observations of Blue and Red kilonova components show that both light ($A \lesssim 120$) and heavy ($A \gtrsim 120$) elements are produced. No direct evidence of individual elements.
  - How can we determine composition?
  - What were the heavier elements produced in the merger?
  - How does the nucleosynthesis depends on merging system?
  - What is the contribution of mergers to light r process elements?

Bibliography:

Cowan, *et al.*, *Making the Heaviest Elements in the Universe: A Review of the Rapid Neutron Capture Process*, arXiv:1901.01410 [astro-ph.HE]